

THE FIRST DETAILED ABUNDANCE ANALYSIS OF GALACTIC BULGE K GIANTS IN BAADE'S WINDOW

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ABSTRACT

We report analysis of echelle spectra ($R = 17,000$, $S/N \approx 50$) of 12 Galactic bulge K giants in Baade's window from the sample of Rich (1988). We perform an abundance analysis for 11 stars ranging from $[Fe/H] = -1$ to $[Fe/H] = 0.45$. The accuracy of our abundance scale is confirmed relative to the disk super-metal-rich stars via the metal-rich giant μ Leonis, and to the Galactic globular clusters using a star in NGC 5927.

We investigate the effects of CN blends on abundance analysis for stars more metal-rich than $[Fe/H] = -0.3$, through spectrum synthesis of the region from 6140 to 7600 Å, including in excess of 8000 ^{12}CN and ^{13}CN lines. We confirm Peterson's assertion that blends of atomic lines with CN vibration-rotation lines can affect abundance results at $R = 17,000$; failure to reject affected Fe lines would have resulted in a spurious 0.3 dex increase in the derived Fe abundances of our most metal-rich stars.

We find that one of the most metal-rich Baade's window stars (BW IV-167) has approximately the same metallicity as the super-metal-rich disk giant μ Leo. The abundance scale of Rich (1988), based on measurement of strong Fe and Mg lines, is ≈ 0.3 dex high but may agree with our results if Rich's $[M/H]$ actually measures $[(Fe + Mg)/H]$. Using our abundances to transform Rich's $[M/H]$ distribution, we find that the mean $[Fe/H]$ is -0.25 for the bulge, slightly less than, but not appreciably different from, the mean $[Fe/H]$ for solar neighborhood G and K giants. Surprisingly, this is not in conflict with the Galactic metallicity gradient found for the solar neighborhood. The extant data on $[Fe/H]$ with Galactocentric radius are completely consistent with a zero gradient within the solar circle.

Infrared flux temperatures for a sample of 34 Baade's window K giants observed by Frogel & Whitford (1987) show that the bulge ($J - K$) versus T_{eff} relation is identical to that of the solar neighborhood. We assert that the observed infrared colors reflect the actual effective temperatures, and that infrared colors affected by molecular blanketing cannot explain why the bulge giant branch is as blue as that of 47 Tuc.

Verification of the ($J - K$) temperature scale, and use of our average $[Fe/H]$ results, suggest that the Frogel & Whitford (1987) infrared H-R diagram requires that the mean mass of Baade's window giants can be *no less* than about $1.1 M_{\odot}$.

Abundances for the α -elements are based on spectrum synthesis of lines due to Mg, Ti, Ca, and Si. We find a puzzling pattern in the trends of element abundance ratios with $[Fe/H]$. Mg and Ti are enhanced by ≈ 0.3 dex relative to solar over almost the full $[Fe/H]$ range. In contrast, Ca and Si closely follow the normal trends for disk giants, with solar ratios occurring above $[Fe/H] \approx -0.2$. The enrichment pattern of bulge giants sets them clearly apart from disk giants of the same $[Fe/H]$. For a subset of the sample we estimate $[O/Fe] \approx 0$ using the $[O I]$ line at 6300 Å; the absence of a carbon abundance indicator requires us to assume a solar C/Fe ratio. We speculate that the high Mg abundances seen in ellipticals (Worthey, Faber, & Gonzalez 1992) may reflect a common enrichment process for bulges and ellipticals. We further propose that the high Ti abundances may largely account for the tendency for Baade's window M giants to have later spectral types than solar neighborhood M giants of the same infrared color (Frogel & Whitford 1987).

Other noteworthy results include enhanced Al abundances similar to those seen in globular cluster giants, and (tentatively) enhancement of the r -process element Eu in the metal-rich bulge stars. The iron-peak and s -process elements appear normal relative to Fe. We find one super-Li-rich giant, BW I-194. We also identify BW I-155 as a spectroscopic double.

This work also presents the following results of general use to investigators in the field: (1) In order to obtain a complete set of gf values for our Fe I lines, we find transformations between numerous $\log gf$ sources. In total we provide a list of 208 atomic gf values, for Fe I and numerous other species. (2) We present a list of uncontamin-

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ated continuum regions appropriate for use in metal-rich stars, as well as atomic lines clear from severe CN blending at a variety of metallicities.

Subject headings: stars: abundances — stars: giants

1. INTRODUCTION

In an era when the abundance scales and even characteristic element ratios of disk, halo, and globular cluster stars are becoming established with increasing certainty, stars of the Galactic bulge have up to now been observed only at low spectral resolution. The image constructed from the *COBE* satellite data (Hauser et al. 1990) strikingly reveals the bulge as a clearly distinct stellar population. The unique mix of evolved stars, ranging from RR Lyrae stars to late M giants and OH/IR stars, has been known since the 1950s (O'Connell 1958). The presence of late M giants and RR Lyrae stars in the same volume is clear evidence for a wide range in metal abundance. In terms of its stellar content, the large numbers of late M giants (Blanco 1965) and their presence suggests that at least part of the population is very metal-rich. The integrated spectrum of bulge fields resembles that of distant elliptical galaxies (Whitford 1978; Rich 1990a), and it is reasonable to expect that studies of the bulge give insights into the formation of elliptical galaxies. As we assemble the description of the stellar population with increasing precision, we recognize that we must have more accurate measurements of the population's age and composition if we are to understand its origin and current evolved stellar content. Indeed, accurate assessment of the age from turnoff photometry depends on composition. Measurement of the range in Fe abundance is no longer sufficient; since overall metallicity depends so critically on the abundances of the CNO elements, detailed composition measurements (element abundance ratios) are becoming increasingly important.

Element abundance ratios potentially give us a glimpse of the earliest formation history. In a widely accepted interpretation, the trend of declining O/Fe with increasing Fe abundance observed in halo stars is thought to reflect the onset of Fe-peak nucleosynthesis in core-deflagration supernovae, which take at least 10^8 yr to develop (cf. Wheeler, Sneden, & Truran 1989). In this paradigm, Fe-peak abundance of a stellar population can be considered to be a kind of time measurement. Matteucci & Brocato (1990) propose that rapidly forming structures such as bulges and ellipticals might have experienced most of their enrichment from massive supernovae in starbursts; consequently, their oxygen and α -capture element abundances would be elevated. From another point of view, systems with a steep age-metallicity relationship will retain a halo-like element ratio imprint to higher Fe abundances, perhaps in fact over the whole range of Fe abundances. Worthey, Faber, & Gonzalez (1992) use low-dispersion spectra to find that ellipticals may be enhanced in Mg in confirmation of this expected trend.

In addition to testing this paradigm (which is not universally accepted; see, for example, the pseudosecondary enrichment models of Edmunds et al. 1991), element ratios in bulge stars may discriminate among formation scenarios currently proposed for the bulge. It is frequently proposed that the gas from which the bulge formed originated elsewhere, perhaps in the halo (Carney, Latham, & Laird 1990; Wyse & Gilmore 1992).

The most metal-poor bulge giants should have element ratios that reflect this formation history. Starburst scenarios present a plausible alternative hypothesis, especially given that starburst-driven winds may eventually form bulges (Heckman, Armus, & Miley 1990; Sofue & Habe 1992). The scenario of rapid enrichment followed by formation of a stellar disk (unstable to bar modes) attractively explains the bulge's high metallicity and narrow range of stellar ages (Rich 1992). The metal-poor stars retain memory of any enrichment of the protobulge gas, while the metal-rich stars preserve the timescale for enrichment relative to the onset of core-deflagration supernovae.

If we can measure element abundance ratios in the Galactic bulge stars, it might help in elucidating how they relate to other metal-rich populations. None of the Galactic metal-rich populations have well-determined compositions, and basic questions (such as possible severe blending of atomic lines with CN) remain the subject of debate. There are four known significant groups of metal-rich stars in the Galaxy, of which the bulge is one. Others include the local "super-metal-rich" (SMR) stars, the high-velocity metal-rich dwarfs (cf. Grenon 1990), and the "disk" system of globular clusters (Armandroff 1989). While the disk globular clusters are metal-rich relative to the globular clusters and halo, in an absolute sense their metallicities of $[\text{Fe}/\text{H}] \approx -0.5$ dex are similar to that of Arcturus, not the SMR stars in the disk, the bulge, or the Grenon sample. Barbuy & Grenon (1990) report $[\text{O}/\text{Fe}] \sim 0.2$ dex for the metal-rich local dwarfs. While it is frequently suggested that these are bulge stars, it must be emphasized that their perigalactica are never less than 3.5 kpc (Grenon 1990), too distant to qualify for bona fide bulge membership unless they are bulge stars expelled by interaction with a bar. Element abundance ratios may provide clear grounds for distinguishing among the various types of metal-rich stars and provide a constraint on the formation history of those stars, much as isotope ratios betray the origin of terrestrial rocks. It is our ultimate goal to explore how abundance patterns at various bulge locations relate to these groups of metal-rich stars, and the halo. Hints of different element abundance ratios are emerging; Barbuy et al.'s (1992) pioneering analysis of star III-17 in NGC 6553 finds an $[\text{M}/\text{H}] \approx -0.2$, yet, in contrast to the SMR dwarfs, $[\text{O}/\text{Fe}] \approx 0$.

Difficulties with field crowding and obscuration prevented quantitative spectroscopy of bulge giants until the advent of digital detectors. The initial group of studies focused on the well-observed vicinity of the globular cluster NGC 6522, the area known as Baade's window at $(l, b) = (1^\circ, -4^\circ)$, approximately 500 pc from the nucleus. Initial studies at low resolution by Whitford & Rich (1983) and more extensively by Rich (1988, hereafter R88) found a wide range in abundances from -1 to 0.8 dex, with $[\text{M}/\text{H}] = 0.3$ in the mean. Geisler & Friel (1992) use the photometric Washington abundance system and confirm R88's abundance scale. The wide abundance range fits closed-box chemical evolution (Rich 1990b) and also explains the presence of late M giants and RR Lyrae stars in the same volume.

It is reasonable to expect that the M giants should be evolved from the K giants and also be metal-rich; indeed, low-resolution spectra also yield a twice solar mean abundance for the bulge M giants (Terndrup, Frogel, & Whitford 1991; Sharples, Walker, & Cropper 1990). At a given $(J - K)$ color, the bulge giants were shown to have stronger TiO bands and thus later spectral type.

Knowledge of fundamental stellar parameters (luminosity, effective temperature, and gravity) is an absolute prerequisite to quantitative abundance determinations. We also address a long-standing problem with the infrared colors and implied effective temperatures of bulge giants. As soon as good infrared color-magnitude diagrams of the bulge were available (Frogel & Whitford 1982), it became clear that despite its high metallicity, the bulge giant branch has the colors of 47 Tuc rather than those of field stars of the expected metallicity. Frogel & Whitford (1982, 1987) initially attributed the blue colors to peculiar blanketing (perhaps an unknown source of molecular opacity). Were this not the case, it would leave high turnoff masses as the only alternative. The problem of the anomalous blue color of the bulge giant branch has played nearly as significant a role motivating this study as the high metal abundances.

If the giant branch is actually blue, one would have to consider the somewhat controversial possibility that the bulge population is in the mean younger than that of the globular clusters (Rich 1992). The blue giant branch locus raised a general concern that it may be impossible to estimate effective temperatures from broad-band colors. This work solidly finds that the observed colors are consistent with the stellar effective temperatures. The stars are indeed bluer than would be expected for globular cluster-aged stars of their metallicity and luminosity.

Terndrup (1988) used the mean $(B - V)$ and $(V - I)$ colors of bulge giant branches to demonstrate an abundance gradient in the bulge as well as to establish an "alternative" abundance scale giving solar metallicity at Baade's window. Given that such a calibration depends on factors such as blanketing, age, etc. (not to mention intrinsic dispersion in these properties), there is little basis for trusting abundances based on this method.

Toward the end of attempting to resolve these questions quantitatively, this paper reports both high-resolution spectroscopy and effective temperatures for the Galactic bulge giants. We have obtained echelle spectra of 12 out of the 88 Baade's window K giants observed by R88. Our abundances are based on the analysis of CCD echelle spectra obtained using the Air Schmidt camera of the CTIO 4 m telescope. The R88 spectra had 5 Å resolution and typical signal-to-noise ratio $S/N \approx 20$. In contrast, our current study uses spectra with a resolution of 0.36 Å and typical $S/N \approx 50$.

Section 2 describes the observations and reductions, while § 3 addresses atmosphere parameters, inclusion of molecules, gf values, and abundance analysis. Section 4 integrates these results into implications for the Galactic abundance gradient and chemical evolution.

2. OBSERVATIONS AND REDUCTIONS

Spectra of 12 Galactic bulge K giants were taken using the CTIO 4 m echelle spectrograph with the Air Schmidt camera on one night in 1990 June and four nights in 1991 May. The

useful spectral coverage obtained was from about 6160 to 7600 Å, although the exact wavelength coverage depended upon the stellar radial velocities and small differences in the position of the echelle spectrum on the face of the CCD. CCD pixels were binned in pairs along the cross-dispersion direction, to reduce the total readout noise. No on-chip binning was performed in the echelle dispersion direction, as this would have resulted in a loss of spectral resolving power. The average FWHM of the Th-Ar comparison lines was measured to be 2.0 pixels, corresponding to a spectrograph resolving power $R \sim 17,000$, approximately constant over the whole spectrum.

We selected 12 Baade's window (BW) bulge K giants, for which R88 had previously estimated overall metallicities from low-resolution, low-S/N spectra acquired at Las Campanas. The stars were selected to span the full range of metallicities found by R88 for BW giants, from $[Fe/H]$ of -1 to 0.8 . While the need to acquire spectra of $S/N \sim 50$ (per extracted pixel) drove us to select relatively bright stars, the high radial velocities give us confidence of their membership in the bulge. Sadler (1992) points out that many stars brighter than $V = 15.5$ may not be bulge members. Further, the spatial depth of the bulge introduces an intrinsic 1σ dispersion of ≈ 0.3 mag in the apparent magnitude. We are confident about membership of the two brightest stars. BW IV-203 is bolometrically the most luminous star in our sample ($M_{bol} = -3.1$) and has $H\alpha$ emission wings consistent with its high intrinsic luminosity; BW I-322 has a radial velocity of -80 km s^{-1} (unlikely for a foreground disk giant), and we find that its element ratios are consistent with those of the faint bulge stars in this study and are not consistent with those of local disk giants.

Table 1 lists Arp (1965) identification numbers for the stars observed, the V magnitude, the S/N obtained, and the total integration time. We also observed μ Leo, a disk SMR star, which we included in our program as a standard for calibration of our abundance scale onto the system of others. In addition to the program star spectra, we also acquired spectra of relatively bright stars for calibration purposes: metal-poor giants with few but narrow lines were observed in order to define the echelle blaze function, and spectra of hot, rapidly rotating B stars were obtained in order to identify telluric features which contaminate lines in the program stars.

TABLE 1
LOG OF OBSERVATIONS

STAR	V	EXPOSURE (minutes)	S/N		
			1990	1991	Total
I-012	15.70	120	...	55	55
I-141	15.83	280	21	57	60
I-155	16.34	160	...	40	40
I-194	16.12	295	26	39	46
I-322	14.21	55	...	50	50
III-152	16.06	167	...	45	45
IV-003	15.06	185	26	60	65
IV-025	16.01	295	19	40	44
IV-072	16.17	150	...	50	50
IV-167	16.53	140	...	30	30
IV-203	13.81	75	...	54	54
IV-329	15.11	120	...	70	70

At all times we obtained at least three frames for each star, which allowed the effects of cosmic-ray strikes on the spectrum to be eliminated. The total integration time for each star varied according to stellar brightness and sky clarity. In 1990 the sky was of particularly low transmission, and the seeing poor; this resulted in noisy sky subtraction because the stellar image affected much of the 10" slit decker. Sky conditions in 1991 were excellent, allowing most of the starlight to be captured by the 2" slit. The stellar images occupied only a small portion of the decker, leaving enough room on the CCD for well-defined sky spectra.

Quartz lamp spectra were acquired during daylight hours immediately before and after the night's observing program. Usually we acquired 21 quartz spectra per night, each of extremely high S/N (in excess of 100 pixel⁻¹ in each frame), in order to define the CCD pixel-to-pixel sensitivity. After the usual trimming and bias subtraction, the quartz frames were grouped into threes, and the median taken to eliminate possible cosmic-ray events. The set of seven median frames were then averaged to form the final quartz frame. The IRAF procedure, APNORM, was then used to convert the final quartz frame into a flat-field frame, containing the pixel-to-pixel CCD pattern noise information.

We extracted the stellar spectra using the IRAF "echelle" package. The echelle format produced 17 orders with ends overlapping in wavelength; we reduced this into a single long spectrum for each star. Aside from the improvements in organization, this resulted in spectra of more uniform S/N. In metal-rich stars, continuum points are scarce, and this format is much superior in establishing the continuum for these difficult cases.

Our initial experiments in joining overlapping echelle orders gave very poor results, because we used old extraction options in the APSUM algorithm. In this algorithm an aperture is selected, over which the light in the order is to be integrated. One usually truncates the apertures to exclude the wings of the order profile; straight inclusion of the light from these wings would actually decrease the final S/N, due to the increasing background noise accompanied by the extra small elements of signal. With spectra extracted in this way we found that the ends of adjacent orders did not match up, and that there was a large-scale ripple (with a wavelength of one order) in the continuum level. These problems occurred because the apertures were not excluding the same fraction of the total signal at different positions along the order, which we believe was caused by the change in the order profile width over the face of the CCD. The change in profile width over the CCD was probably due to small deviations from perfect focus at different positions.

To overcome the problem we used the new IRAF extraction algorithm APALL, which performs "optimal extraction" (see Horne 1986). This algorithm permitted use of wide extraction windows, including all flux from the order, without degrading the S/N of the spectra.

After bias subtraction, trimming, and flat-fielding, the optimal extraction procedure was used to extract two-dimensional echelle frames from each CCD image. Since at least three images were taken per star, it was then possible to remove the effects of cosmic-ray strikes. The cosmic-ray removal algorithm first scales the frames by the mode of the counts over the

whole extracted image; then for each pixel in the spectrum the median is adopted as representative of the true spectrum flux. Data values for each pixel in each frame are examined, and pixels where the flux is farther than 3 σ from the median are rejected. Since cosmic rays affect more than 1 pixel, we ignore data within 3 pixels of each rejected point, thus eliminating all possible effects of cosmic rays from the averaged spectrum; pixels not rejected are then averaged to form the adopted flux for each point in the spectrum.

The average two-dimensional echelle frame was wavelength-calibrated using the Th-Ar spectra. The IRAF routine DISPCOR was used to sum overlapping portions of orders, producing a single spectrum, this spectrum being modulated by the echelle blaze function. The form of the blaze function was found by fitting the continuum of the bright metal-poor K giant star with a high-order polynomial (with IRAF routine ECCONTINUUM). The continuum fit was then wavelength-calibrated to produce a single smooth spectrum, modulated by the blaze function. The average, wavelength-calibrated spectrum was then divided by the blaze function spectrum to produce the final result. To test the validity of the method, we repeated the process, but without co-adding overlapping orders; this enabled us to overplot the portions of adjacent orders. We found these overlapping portions of spectrum to be in excellent agreement, and consistent with the S/N of the spectra; this gave us confidence in the method.

The S/N values given in Table 1 were obtained by measuring the standard deviation of the ratio of two individual extracted spectra; the noise in this ratio is the quadrature sum of the noise from the individual spectra. In this way it was possible to estimate the noise on the sum of the three spectra.

3. ANALYSIS

Before a quantitative analysis the spectra were inspected for unusual characteristics. In Figure 1 we show a portion of spectrum for the star BW I-155, which is particularly complex due to a doubling of the absorption lines. The double hydrogen line feature in Figure 1 is clearly visible, as well as a doubling of numerous atomic lines. The cores of the H α pair in Figure 1 are much shallower than Balmer lines in single stars, which is consistent with a double K giant system; the radial velocity separation is approximately 80 km s⁻¹. Note that the spectra

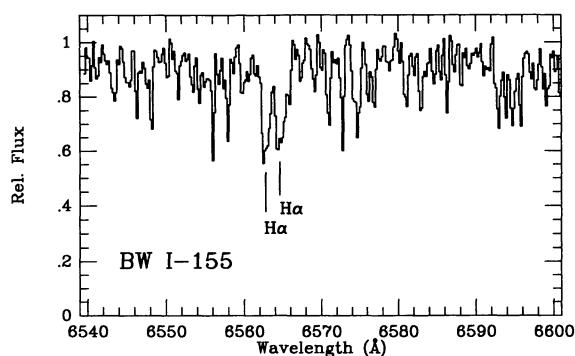


FIG. 1.—Spectrum of BW I-155 showing double H α line, separated by approximately 80 km s⁻¹. Other line pairs are clearly visible with nearly equal strength. Notice the greatly reduced H α central depth.

in Figure 1 are of similar line strength, which indicates that the stars are of similar spectral type and metallicity.

The single observation of a radial velocity separation suggests that BW I-155 is either a spectroscopic binary or a line-of-sight double system. The observed radial velocity separation of the components of BW I-155 suggests that if it is a binary system, then the orbital period lies approximately in the range 20–60 days, with a semimajor axis between 70 and 180 $R_{\odot}$. Since BW I-155 is slightly fainter than BW IV-167, it is reasonable to suppose that BW I-155 is of lower luminosity, with $\log g$ near 2.0 (see Table 7) and $R \sim 17 R_{\odot}$, in which case both stars can easily fit inside the dimensions of the implied orbit. If the gravities of the two K giants are closer to $\log g \sim 1.0$, then the large stellar radii exclude the possibility of a binary system. Additional spectroscopic observations are required to resolve this issue. If BW I-155 is indeed a spectroscopic binary, future observations might yield masses for the components, and therefore an estimate for the bulge turnoff mass.

In Figure 2 we show a portion of spectrum around $H\alpha$ in BW IV-203, indicating the presence of emission in the blue and red $H\alpha$ wings. This is of considerable importance, since BW IV-203 is our most intrinsically luminous star (at $M_{\text{bol}} = -3.1$) and would in principle be a candidate for a foreground disk giant. Figure 2 also illustrates the quality of our spectra; in particular, the well-defined line of Ti II at 6559 Å, with an equivalent width (EW) of 55 mÅ, implies that lines of only 20 mÅ can readily be detected.

A rare Li-rich giant also occurs in our sample. In Figure 3 we show data for four stars in the Li I region at 6708 Å. The Li I line EWs in our bulge giants are similar to those seen in disk giants (at ~ 30 mÅ), with the exception of BW I-194, which possesses a Li line EW in excess of 200 mÅ. The strong Li line in BW I-194 puts this star in the class of super-lithium-rich giants; the frequency of these stars among solar neighborhood giants is approximately 1% (Brown et al. 1989).

3.1. Equivalent Width Measurement

Equivalent widths of useful lines were measured with a semi-automated program, GETJOB, an improved version of GETPHD, described by McWilliam (1988, 1990), but with improved continuum fitting routines. Unlike GETPHD, it

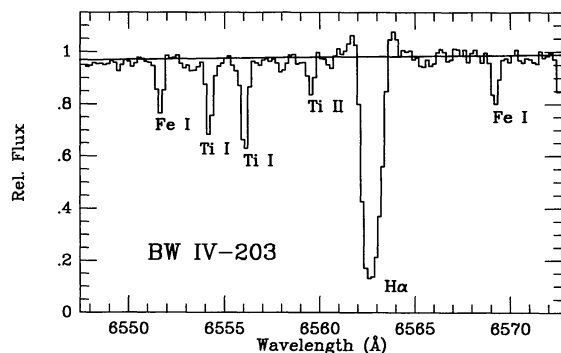


FIG. 2.—Spectrum of BW IV-203 near $H\alpha$, showing emission components in blue and red $H\alpha$ wings. Solid horizontal line represents the local continuum. The 55 mÅ Ti II line near 6559 Å demonstrates that line equivalent widths of 20 mÅ can readily be measured.

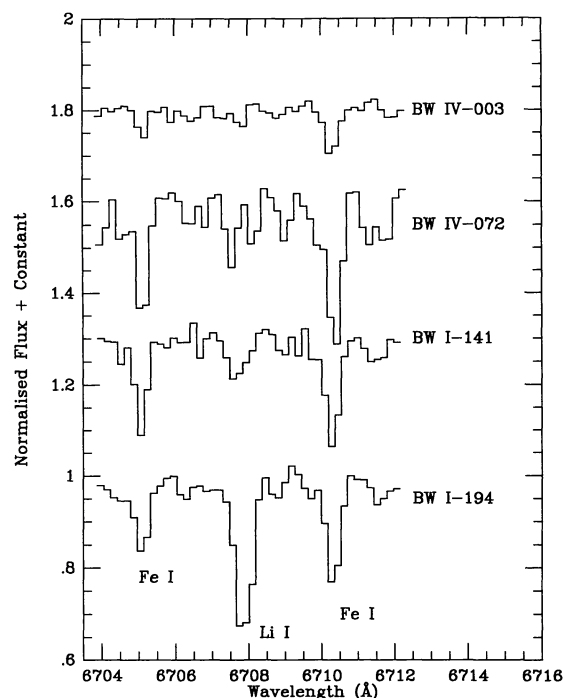


FIG. 3.—Spectra of Baade's window giants near the Li I line at 6708 Å. A Li I equivalent width of 213 mÅ for BW I-194 puts this star in the class of super-Li-rich giants.

does not require a cross-correlation template spectrum in order to find the stellar radial velocity; rather, it relies on $H\alpha$ and a list of strong, clean, atomic lines. A full account of the operation of GETJOB is not warranted here, but the method of continuum placement is noteworthy. In the mode used for all program stars except BW IV-167 (due to low S/N) a semiautomated continuum placement method was adopted. The routine finds high points of contiguous pixels within a set of windows; often the highest 6 pixels were found in each window of 200 pixels. Then a polynomial fit is made to the continuum points found over the entire spectrum. Such methods of finding high points are prone to systematic errors, due to the bias of using statistically high pixel groupings. In GETJOB, however, this problem has been addressed by computing the probable shift from the true continuum level, based on the S/N of the spectrum and the amount of line blanketing. GETJOB then adopts a continuum level for the region based on the computed shift. Obviously, selecting a few pixels in every 200 in the SMR star μ Leo is unlikely to pick out statistically deviant points because the spectrum is very heavily blanketed. In practice, if the shift turned out to be large, then the number of contiguous pixels, or the search window size, was altered, to minimize the magnitude of the shift. Typical shifts turned out to be of order 1% or less of the continuum flux level.

The polynomial fits to the continuum regions over the entire spectrum utilize a χ^2 method to find the order of the best fit, are inspected by eye, and when necessary are manually overridden. Typically, the scatter about the continuum fit was 1%–2% of the continuum level.

For the spectrum of the SMR bulge star BW IV-167 (R88) the S/N is low and the blanketing high, making continuum

placement by the above method somewhat uncertain. Therefore, for this star only the continuum regions found for μ Leo were used, since the stars are of similar metallicity. In Table 2 we list the wavelength bounds of the μ Leo continuum regions which should be of use to those wishing to study SMR stars in general. The validity of these continuum regions was checked by inspecting our μ Leo spectrum and the α Boo spectrum of Griffin (1968), but no obvious contaminating features were found. This method, for BW IV-167 continuum placement, requires no estimate of continuum shifts because no statistical selection effects were involved.

Comprehensive details of the operation of GETJOB will be published in a forthcoming paper; we note, however, that the interactive display feature of GETJOB allows the user to visually inspect, and if necessary to correct, the nonlinear least-squares Gaussian line fits and polynomial fits to the continuum made by the program. This ensures that gross errors are unlikely to survive the measurement process. Equivalent widths measured with these routines have been shown (McWilliam 1988) to be in good agreement with the results of other workers; the results are particularly reliable for spectra of high S/N, containing plentiful continuum.

Line identifications and wavelengths were mostly taken from Moore, Minnaert, & Houtgast (1966); we searched for clean lines at a resolving power of $R = 17,000$. Particular attention was paid to the α -element lines of Mg, Si, and Ca, as well as numerous heavy-element lines and lines from Na and Al. Although we tried to ensure that these lines were clean in the solar spectrum, the same cannot be said for the lines measured in the SMR stars, where blends are much more frequent.

For perfect removal of telluric absorption from our spectra it would have been necessary to acquire spectra of hot stars immediately prior to and after each program star spectrum. However, in order to maximize time spent on our program stars, only one hot star spectrum was obtained on each night. We were unable to scale these spectra in a manner which would allow for complete division of all telluric absorption. Therefore, lines affected by telluric absorption were identified by overplotting the telluric and program star spectra. When contaminated lines were found, a simple correction was applied to the line EW if the telluric contamination amounted to less than ~ 20 mÅ. In cases where the telluric EW was greater than

about 20 mÅ, or the telluric line was larger than $\sim 20\%$ of the target line EW, the stellar line was not used.

In Table 3 we list EWs of 2287 lines measured by GETJOB for 11 Baade's window giants and μ Leo; 208 EWs for α Boo, hand-measured from the atlas of Griffin (1968), are also listed. We include α Boo in order to detect zero-point problems, such as gf values or non-LTE effects on individual lines, and as a standard to compare with the analyses of other workers. Line EWs in Table 3 with the superscript "b" have been corrected for telluric absorption; lines with excessive telluric absorption have been eliminated from Table 3. Note that EWs have not been measured for BW I-155, owing to the complex nature of its double-lined spectrum.

3.2. Atmosphere Parameters

In order to compute accurate abundances from EWs, it is necessary to have appropriate parameters which define the model atmosphere to be used in the analysis. The parameters needed are effective temperature (T_{eff}), gravity ($\log g$), metallicity ($[\text{Fe}/\text{H}]$), and microturbulent velocity (ξ). The effective temperatures can be accurately determined from broad-band photometry, provided that the reddening is known. For GK giants ($V - K$) is a particularly good indicator of T_{eff} , because the index changes rapidly with temperature, making it relatively insensitive to errors. The ($J - K$) index is also useful because of the insensitivity to line blanketing, which may occur in the metal-rich stars. The ($J - K$) index is usually inferior to ($V - K$) because ($J - K$) is less sensitive to temperature than ($V - K$) in GK giant stars; but it is superior to the ($R - I$) index due to the diminished reddening at infrared wavelengths.

In Table 4 we review the current IR photometry, V dereddened ($V - K$) and ($J - K$) indices, and M_{bol} based on a reddening of $E(B - V) = 0.48$, and a distance of 8 kpc. We use the reddening and color corrections adopted by Frogel & Whitford (1987, hereafter FW87). This reddening is in good agreement with the value found for BW RR Lyrae stars by Walker & Terndrup (1991), $E(B - V) = 0.50 \pm 0.05$.

Typical uncertainties about the mean reddening value are usually estimated at $\Delta E(B - V) = \pm 0.05$ mag.

In Table 5 we show the effect of a change in $E(B - V)$ of 0.05 mag on the computed effective temperature for ($V - K$), ($V - J$), and ($J - K$) indices; inspection of the table shows that the ($V - K$) color is superior. The temperatures in Table 5 were computed using the McWilliam (1990, hereafter McW90) calibration of broad-band colors with T_{eff} . The temperature uncertainties listed in Table 5, resulting from $E(B - V)$ error, are used as weights for determining the mean temperatures, because it is expected that the reddening uncertainty is the chief source of error in T_{eff} . Using the McW90 calibrations, the weighted mean effective temperatures have been computed for all program stars and are listed in Table 6.

Inspection of Table 6 shows that T_{eff} derived from ($J - K$) appears to be greater than ($V - K$)-based temperatures for the R88 SMR bulge stars in our sample. In Figure 4 we present R88 $[\text{Fe}/\text{H}]$ against ΔT , where ΔT is $T(J - K) - T(V - K)$. Figure 4 shows that the average ΔT is close to zero below solar metallicity, but for the SMR bulge stars the ($J - K$) temperature is larger than the ($V - K$) temperature by 250–300 K.

We suggest that this temperature difference is due to blan-

TABLE 2
CONTINUUM REGIONS IN μ LEO (Å)

Blue	Red	Blue	Red
6205.64	6206.09	6908.99	6909.44
6283.47	6283.92	6995.63	6996.08
6325.53	6325.98	7047.75	7048.20
6372.08	6372.53	7075.08	7075.53
6446.68	6447.13	7157.04	7157.49
6467.71	6468.16	7168.01	7168.46
6539.79	6540.24	7241.52	7241.97
6577.90	6578.35	7296.34	7296.79
6626.61	6627.06	7373.46	7373.91
6723.85	6724.30	7404.37	7404.82
6758.00	6758.45	7434.03	7434.48
6787.84	6788.29	7556.80	7557.25
6862.97	6863.42		

TABLE 3
MEASURED EQUIVALENT WIDTHS

ID	λ Å	1012	1141	1194	1322	3152	4003	4025	4072	4167	4203	4329	α Boo	μ Leo
Li I	6707.800	20	40	215	15 ^a	15 ^a	13	15 ^a	15 ^a	25 ^a	23	13	3 ^a	28
Na I	6154.230	...	84	118	108	62	...	125	168	152	...	31	69	170
Na I	6160.753	111	162	103	157	102	...	152	...	172	59	48	...	198
Mg I	6319.242	80	108	166	104	73	...	97	111	157	66	44	52	127
Mg I	6894.890	...	80	88	...	47	16	68	...	180	23	...	48	...
Mg I	6965.408	...	66 ^b	...	51 ^b	62	...	90	131	153 ^b	32 ^b	28 ^b	51	118
Mg I	7387.700	82	144	112	112	83	40 ^b	150	128	142	40	63	83	160
Al I	6696.032	149	140	161 ^b	171	150	29	171	207	212	92	66	92	193
Al I	6698.669	95	102	90	87	91	...	117	116	127	40	62	63	104
Si I	6237.328	63	81	77	82	45	...	75	170	120	54	36	63	137
Si I	6721.844	52	35	41	62	72	26	81	83	94	21	24	43	93
Si I	6976.504	49	67	47	60	89	...	73	99	122	19	29	32	...
Si I	7226.208	43	68	35	...	60 ^b	28	59	...	125	27	...	33	...
Si I	7235.325	35	...	28 ^b	...	...	...	31 ^b	...	...	...	...	28	...
Si I	7405.790	106	110 ^b	80	92 ^b	95	54	98	110	112 ^b	65 ^b	84	80	147
Si I	7415.958	94	115 ^b	88	120	110	65	160	131	199	36	63	78	184
Si I	7423.509	179	231	151	210	92 ^b	87	255	...	...	98	75	98	313
Sc I	6210.671	133	116	186	143	149	...	114	177	153	154	58	116	143
Sc II	6604.600	95	106	86	131	108	65	96	99	94	91	87	87	109
Ca I	6166.440	157	149	133	139	95	83	128	143	150	121	125	121	134
Ca I	6169.044	145	154	217	201	192	134	186	189	207	148	140	138	182
Ca I	6169.564	206	165	181	164	172	95	192	132	197	189	150	177	197
Ca I	6417.685	59	71	81	86	66	23	95	115	116	23	20	50	98
Ca I	6455.605	120	132	131	159	145	86	133	136	167	131	89	109	135
Ca I	6464.679	84	82 ^b	90	106	82	...	107	125	122	70	43	71	104
Ca I	6508.846	56	45 ^b	67	88	50	...	88	126	144	41	35	44	103
Ca I	6572.795	195	160 ^b	226	250 ^b	192	85	177	224	225 ^b	216 ^b	133	171	214
Ca I	7148.150	219	221 ^b	242	228 ^b	228	149	238	285	255 ^b	242 ^b	177	204	256
Ca I	7326.160	213	201	165 ^b	235 ^b	...	138	261	332	308	170	130 ^b	166	...
Ca I	6719.687	32	26	24	30	...	...	...	...	90	18	20	17	64
Ti I	6303.767	109	106	156	149 ^b	113	33	77	131	160 ^b	117 ^b	67 ^b	88	173
Ti I	6336.113	95	76	114	105	78	22	117	137	117	95	43	82	146
Ti I	6554.238	129	132	148	147 ^b	112 ^b	63 ^b	136	153	170	129	79 ^b	107	150
Ti I	6556.077	164	138	159	169	132 ^b	58	152	181	191	167	113	142	158
Ti I	6743.127	161	140	168	186	188	63	141	192	182	182	119	143	155
Ti I	6861.500	86	111	112	88	127	21	85	...	137	...	53	64	133
Ti I	7209.468	195	...	211	204	212	91	...	...	218	209	136	173	216
Ti I	7216.190	...	145	160	184	153	48	...	180	203	162	105	119	194
Ti I	7251.717	...	149	...	187 ^b	133	86	142	...	201 ^b	183 ^b	131	133	...
Ti II	6559.576	78	72	73	88	40 ^b	74	79	49 ^b	72	55	48	63	87
Ti II	7214.741	85	77	70	108	78	45	116	133	141	86	48	48	186
V I	6274.658	155	114	150	154	146	...	111	189	173	172	98	130	145
V I	6285.165	139	179	140 ^b	242	...	...	118	213 ^b	216	236 ^b	95 ^b	122	121
V I	6504.186	81	70	80	105	78 ^b	10	106	118	163	61	44	62	110
V I	6531.429	109	106	104	145	122 ^b	10 ^b	105	141	137 ^b	91	79	88	120
V I	6543.510	48	79	119	91	66 ^b	...	82 ^b	152	127 ^b	28 ^b	27	36	110
V I	6812.356	36	30	56	58	53	...	54	98	106 ^b	29 ^b	23	35	83
Cr I	6330.096	124	117	137	141	127	51	109	122	155	138	83	100	134
Cr I	6881.716	...	81	...	...	...	...	94	...	135	42	...	60	136
Cr I	7400.188	153	178	145	184	169	89	170	193	209	159	126	121	193
Cr I	7462.342	169	169	179	174	171	100	200	200	226	155	129	146	206
Fe I	6151.623	...	97	85	124	94	...	...	...	150	...	89	105	127
Fe I	6173.341	147	181	148	180	178	101	168	173	207	153	92	145	168
Fe I	6180.209	160	177	125	169	157	93	163	185	189	123	115	115	186
Fe I	6187.995	79	80	85	98	92	53	101	102	95	81	64	80	108
Fe I	6200.321	150	156	147	153	156	96	122	171	188	155	124	124	167
Fe I	6212.067	55	42	36	56	38	...	60	91	95	...	...	15	87
Fe I	6213.437	162	194	198	236	123	115	168	133	191	208	130	148	263
Fe I	6219.287	182	196	179	210	175	122	173	190	251	209	140	171	215
Fe I	6229.232	102	119	94	138	83	51	104	123	149	118	70	93	130

TABLE 3—Continued

ID	λ Å	1012	1141	1194	1322	3152	4003	4025	4072	4167	4203	4329	α Boo	μ Leo
Fe I	6240.653	122	109	92	115	67	70	112	...	166	111	64	121	172
Fe I	6246.327	169	160	160	173	165	120	187	...	231	162	124	139	184
Fe I	6254.253	210	216	198	223	196	129	198	193	224	196	162	184	235
Fe I	6265.141	201	200	193	246	186	139	223	244	246	213	151	190	260
Fe I	6270.231	115	113	109	122	103	75	101	124	152	142	98	93	133
Fe I	6271.283	78	75	80	92	81	...	81	81	75	58	40	61	117
Fe I	6297.799	165	176	173	187 ^b	168	136	170	178 ^b	225	193	...	140	...
Fe I	6301.508	...	165	142	166 ^b	80 ^b	111 ^b	203	160 ^b	218	152	114	150	...
Fe I	6307.854	...	...	...	...	17	12 ^b	42	37	82	30	...	18	54
Fe I	6311.504	101	131	101	182	66	47 ^b	99	126	141 ^b	98 ^b	54 ^b	81	194
Fe I	6322.694	144	147	148	171	157 ^b	101	143	179	203	155	113	127	185
Fe I	6330.852	58	29	44	80	50	15	75	62	74	15	37	40	92
Fe I	6335.337	206	195	177	136	178	130	213	251	231	201	157	187	223
Fe I	6336.830	150	148	147	159	123	101	158	177	183	140	124	134	189
Fe I	6353.849	56	56	63	101	45	...	80	80	146	59	22	35	100
Fe I	6355.035	165	173	209	216	152	82	186	232	261	150	132	161	212
Fe I	6380.750	85	122 ^b	100	140	90	39	131	102	196 ^b	68	61	86	161
Fe I	6392.538	79	91	94	149	60	23	106	113	110	80	60	62	111
Fe I	6393.612	236	251	289	286	221	169	239	312	322	230	213	216	279
Fe I	6411.113	...	36	47	43 ^b	36	...	...	71	...	...	7	14	79
Fe I	6411.658	157	171	166	189	133	106	207	202	226	159	138	156	182
Fe I	6419.956	133	142	156	154	140	58	179	175	148	92	88	98	175
Fe I	6421.360	225	225	212	244	177	139	214	236	235	206	172	202	240
Fe I	6430.856	251	241	265	272 ^b	226	152	257	316	304	244	188	176	280
Fe I	6475.632	129	...	121	149	136	83 ^b	128 ^b	139 ^b	114 ^b	119 ^b	96	136	153
Fe I	6481.878	132	145	153	176 ^b	...	95	159	156 ^b	168	172	111 ^b	139	170
Fe I	6494.994	292	267	261 ^b	333 ^b	269	165	369	375	376	269	218 ^b	270	...
Fe I	6498.945	166	165	150	211	162	101	152	198	215	182	137	157	186
Fe I	6518.373	123	121	153	...	112 ^b	83	138	156 ^b	154	117	...	119	148
Fe I	6533.940	69	91	67	100 ^b	67	30	121 ^b	101	99	43	...	60	134
Fe I	6546.252	189	196	195	204	185	136	192	204 ^b	196	196	165	166	209
Fe I	6556.806	17	27	...	35	18	...	57	35	52	12	12	23	45
Fe I	6569.224	116	106	128	168	101	58	158	158	171	83	92	94	163
Fe I	6571.180	...	33	27	57	...	...	36	60	61	9	...	15	...
Fe I	6574.254	138	114	95	157 ^b	134	68	108	123	135	142	101 ^b	114	128
Fe I	6575.037	140	131 ^b	148	187	147	85	170 ^b	220	209	125	106	119	204
Fe I	6591.306	26	29	47	52	24	...	45	36	62	...	16	16	55
Fe I	6593.874	163	171	148	225	161	100	174	200	225	151	148	182	196
Fe I	6597.571	58	72	64	136	64	40	95	95	89	30	42	48	86
Fe I	6608.044	97	100	103	149	85	35	108	133	133	80	76	62	110
Fe I	6625.039	195	177	197	221	201	63	173	229	233	169	116	92	218
Fe I	6627.560	59	54	39	66	85	27	82	70	102	26	35	44	81
Fe I	6646.966	76	76	68	101	47	22	86	92	119	47	46	44	103
Fe I	6648.121	110	88	71	123	89	27	106	142	145	90	49	72	141
Fe I	6653.911	31	27	41	39	...	19	52	40	71	15	25	28	64
Fe I	6677.997	220	227	220	254	232	148	260	275	288	187	184	202	278
Fe I	6699.136	19	...	21	27	26	20	...	37	20	...	13	16	50
Fe I	6703.576	111	109	117	137	124	59	139	135	119	98	84	80	134
Fe I	6704.500	27	...	19	34	...	...	39	31	29	...	16	15	45
Fe I	6705.105	85	81	66	93	85	30	108	108	72	45	44	56	108
Fe I	6710.323	112	101	108	137	108	60	136	154	168	120	75	87	153
Fe I	6713.745	39	42	34	55 ^b	25	...	65	53	96	17	15	31	76
Fe I	6715.386	66	70	59	98	61	...	90	111	125	28	29	41	154
Fe I	6716.252	70	...	71	76	...	15	79	79	94	...	10	32	131
Fe I	6725.364	35	45	35	62	47	...	73	53	71 ^b	25 ^b	30	38	71
Fe I	6726.673	64	64	65	87	61	27	74	96	134	28	48	62	91
Fe I	6733.153	57	49	43	61	44	17	70	65	73	22	25	42	67
Fe I	6739.524	95	78	76	109	68	...	87	116	102	94	63	75	115
Fe I	6750.164	149	146	138	154	151	97	130	154	154	158	133	146	156
Fe I	6752.716	200	140	217	246	100	24	206	358	292	58	54	53	260
Fe I	6806.856	107	94	92	111	110	48	114	120	134	85	77	92	135
Fe I	6810.267	81	75	67	94	83	22	100	112	115	53	54	66	95
Fe I	6820.374	72	62	85	81	70	28	82	95	101	49	46	69	102
Fe I	6828.596	83	89	94	102	92	36	108	116	117	57	66	73	126
Fe I	6839.835	105	103	114	130	108	40	139	153	161	94	73	88	158
Fe I	6841.341	151	152	163	161	130	45	193	213	203	68	83	82	215
Fe I	6842.689	72	88	72	83	66	31	78	73	85	40	36	60	105
Fe I	6843.655	93	112	86	110	99	54	113	120	141	70	73	87	123
Fe I	6851.652	76	87	99	88	90	...	87	88 ^b	121	57	45	51	108
Fe I	6855.723	53	69	42	70	49	28	80	...	97	30	29	41	123
Fe I	6858.155	85	113	82	78	87	48	91	101	122	59	54	66	130
Fe I	6898.307	...	85	66	...	36	13	49	64	124	29 ^b	...	38	100

TABLE 3—Continued

ID	λ Å	1012	1141	1194	1322	3152	4003	4025	4072	4167	4203	4329	α Boo	μ Leo
Fe I	6911.522	...	86	88	...	...	17 ^b	77	...	92	57	...	62	114
Fe I	6916.686	104	93	92	149	...	39	112	...	128	74	70	79	...
Fe I	6960.330	39	44	64	61	...	...	105	102	...	...	17	25	...
Fe I	6978.862	162	163	212	228	148	130	167	150	215	198	144 ^b	129	295
Fe I	6988.533	98	92 ^b	86	...	82 ^b	78	...	90	105 ^b	85 ^b	...	88	102
Fe I	7007.976	72	59	87	69 ^b	62	22	70	...	121	39	30	56	114
Fe I	7022.957	...	115 ^b	101	99 ^b	92 ^b	...	90 ^b	103	128	71 ^b	83	85	127
Fe I	7024.644	91	125	102	172 ^b	80 ^b	22 ^b	162 ^b	215	137 ^b	...	52	67	243
Fe I	7038.220	102	158 ^b	102	160	89	48	207	...	...	...	76	76	242
Fe I	7068.423	124	167	147	153	121	75	189	197 ^b	190	116	86	107	232
Fe I	7071.866	47	73	30	45	41	...	71	60	68	28	29	40	75
Fe I	7083.394	...	44	44	68	28	...	48	50	...	...	1	34	124
Fe I	7090.390	129	125	129	...	125	...	154	193	157	89 ^b	53 ^b	87	212
Fe I	7112.170	120	102	106	131	113	33	132	150 ^b	140	76	51 ^b	77	167
Fe I	7127.573	75	98	73	113	43	21	87	139	159	18	27	36	159
Fe I	7130.925	142	155	178	179	109	82	209	269	301	122	111	114	241
Fe I	7132.985	85	99	117	122	102	30 ^b	80	156	184	46	60	65	157
Fe I	7142.517	70	86	111	92 ^b	72	21	117 ^b	172	147	36	47	50	163
Fe I	7145.312	109	94	100 ^b	122 ^b	66 ^b	30	155	210	212	59 ^b	40 ^b	69	221
Fe I	7151.464	103	100 ^b	121	127	113 ^b	46	112 ^b	149	161	91	61 ^b	73	146
Fe I	7155.634	101	148	118 ^b	148	70 ^b	...	144	288	...	52	31	44	168
Fe I	7180.004	...	103	93	...	...	111	...	177 ^b	111 ^b	67	101	...	...
Fe I	7219.680	77	91	73	87	73 ^b	...	83	...	104	68	49 ^b	77	114
Fe I	7228.700	90	111	...	138	104 ^b	...	141	196	...	...	63	77	182
Fe I	7256.142	...	17	24	...	...	...	35	...	67	...	...	23	...
Fe I	7288.740	...	...	116	139	113 ^b	...	163	230	...	...	...	105	...
Fe I	7396.526	21	...	...	29	28	...	57	28	33	...	23	33	71
Fe I	7411.162	127	152 ^b	137	176	150	100	184	189	196	127	112	119	208
Fe I	7418.672	86	125	88	105	64 ^b	48	99	94 ^b	153	73	65	68	135
Fe I	7420.241	19	31	...	21	...	21	35	33	87 ^b	...	...	14	82
Fe I	7421.560	32	...	42	35	43	...	35	37	68	7	13	35	72
Fe I	7445.758	184	176	176	189	192	105	250	242	293	155 ^b	136	130	256
Fe I	7454.004	47	79	30	47	44	...	58	65	...	9	19	34	96
Fe I	7461.527	118	123	113	114	90	36	98	163	...	81	67	71	154
Fe I	7477.595	34	37	36	70	31	...	64	62	85	16	13	30	73
Fe I	7491.652	128	116	156	153	63	58	101	144	160	97	63	76	149
Fe I	7498.535	68	62	83	82	58	...	56	100	70	...	31	42	117
Fe I	7506.030	50	60	47	95 ^b	30	25	107	86	110	20	8	29	124
Fe I	7507.273	93	109	124 ^b	122	89	43 ^b	139	136 ^b	128	75	68	74	155
Fe I	7531.153	123	132	121	153	138	91	169	168	194	120	101	94	183
Fe I	7540.444	64	99	64	66	75	29	80	88	91	48	25	48	102
Fe II	6416.928	26	20	60	68	39	23	48	55	53	29	36	39	52
Fe II	6432.683	53	74	42	61	40	50	63	65	63	33	49	41	64
Fe II	7449.338	36	54	31	39	26	15	54	70	89	24	17	22	104
Co I	6188.998	131	165	135	150	101	44	120	144	137	82	98	104	190
Co I	6678.849	75	72	104	91	56	...	70	157	130	38 ^b	35	50	133
Co I	6814.961	136	126	131	130	139	49	126	177	184	102	88	121	173
Co I	7052.870	181	...	175	161 ^b	175	62 ^b	184 ^b	213	236	149	110 ^b	188	232
Co I	7417.390	140	168	111	156 ^b	90	31	107	177	170	61	73	100	130
Ni I	6327.604	128	119	117	144	110	68	96	164	126	121	102	109	153
Ni I	6339.118	109	146	122	179	102	60	194	214	170	78	56	66	247
Ni I	6350.495	30	30	38	41	20	...	53	69	88	11	16	17	85
Ni I	6366.491	144	133	173	176	144	27	159	198	225	115	83	120	195
Ni I	6482.809	117	115	121	141	104 ^b	69 ^b	125 ^b	122 ^b	147	107	91	118	143
Ni I	6532.881	105	112 ^b	92	134	103	46 ^b	121	163	150 ^b	76 ^b	58	74	152
Ni I	6586.319	111	113	97	155	107	64	112 ^b	100	139	105	90	97	122
Ni I	6643.638	201	187	170	213	181	132	202	218	233	201	157	184	202
Ni I	6767.784	161	157	154	173 ^b	159	119	127	167 ^b	187	179	159	171	171
Ni I	6772.321	90	92	82 ^b	91	85	43	90	115	139	75	66	72	127
Ni I	7110.905	126	150	120 ^b	154	131	54	143	167	235	101	74 ^b	115	199
Ni I	7122.206	174	182	148	211	143	99	187	230	243	125	145	134	268
Ni I	7291.499	142	...	202	169	148	90 ^b	281 ^b	...	...	...	91	153	...
Ni I	7393.609	139	174	130	156 ^b	153	95	143	169 ^b	209	117	105 ^b	117	219
Ni I	7414.514	173	203	145	184	185	82	176	213	220	164	121	147	207
Ni I	7422.286	131	439	134	144	135	96	169	173	179	124	112	120	190
Ni I	7522.778	157	164	162	196	157	88	173	199	240	117	72	116	206
Ni I	7525.118	113	130	117 ^b	124	140	71	133	139	171	99	87	82	174
Ni I	7555.607	130	139	130	128	101	103	142	162	164	106	110	114	147

TABLE 3—Continued

ID	λ Å	1012	1141	1194	1322	3152	4003	4025	4072	4167	4203	4329	α Boo	μ Leo
Y I	6435.049	49	58	75	81	54	27	81	125	94	67	41	34	93
Y II	6795.410	15	20	26	20	37	21	25	24	51	30	23	15	32
Y II	7450.320	33	50 ^b	...	44	50	17	59	67	117	29	16	17	95
Zr I	6762.398	32	20	53	49	62	26	67	218	86	79	31	21	75
Ba II	6141.727	...	57	...	...	...	...	...	...	...	...	...	161	...
Ba II	6496.908	208	185	158 ^b	258	194	158	172	230	207 ^b	283 ^b	190	173	209
La II	6390.480	35	30	57	50	33	23	29	59	53	46	30	28	61
La II	6774.260	31	25	33	55	34	21	28	31	39	56	32	27	37
Eu II	6645.127	54	48	46	72	50	32	75	73	74	30	38	39	58

^a Upper limit to equivalent width only.^b Corrections made for telluric contamination to equivalent width.

ketting of the V band associated with the high metallicity of these stars. For IV-167 a change in V magnitude of about 0.35 mag would be necessary to account for the observed ΔT .

Given the discrepancy between $(V-K)$ and $(J-K)$ temperatures for the SMR stars in our sample (using $[\text{Fe}/\text{H}]$ suggested by R88), we adopt T_{eff} derived from the $(J-K)$ color only. Adopted effective temperatures and other atmospheric parameters are listed in Table 7.

The reliability of $(V-K)$ and $(J-K)$ colors as temperature indicators for BW giants has been called into question by FW87, who suggested that the supposed high metallicity of

bulge giants results in blanketing of the photometric bands and significantly alters the temperature-color relations. We can test the accuracy of the color temperatures using the bolometric correction data measured by FW87 and the infrared flux method. The infrared flux method is widely considered to be the most accurate fundamental temperature indicator known (Blackwell & Shallis 1977; Blackwell, Petford, & Shallis 1980) and has been used to calibrate the temperature-color relations of solar neighborhood GK giants (Bell & Gustafsson 1989; McW90). Briefly, the method utilizes the ratio R of total flux to flux per unit frequency at an infrared wavelength. This ratio is approximately proportional to the cube of the stellar effective temperature. Values of R computed using LTE model atmospheres are incredibly insensitive to the other model atmosphere parameters (e.g., $\log g$ and $[\text{Fe}/\text{H}]$), and even to different grids of model atmospheres taken from the literature. The uncertainty in T_{eff} found from the infrared flux method is approximately one-third of the uncertainty in the measured fluxes (McW90); claims of 6% accuracy in absolute fluxes (Leggett et al. 1986) suggest that T_{eff} can be measured with an accuracy of $\sim 2\%$, or about 90 K in K giants. Our experiments with numerous model atmospheres, including outdated atmosphere grids, show that the contribution to the temperature uncertainty arising from the inaccuracies in the model atmospheres amounts to approximately ± 30 K.

It is clear that the ratio R is intimately related to the infrared K -band bolometric correction (BC). We present equation (1), which allows IR flux method R -values to be computed from BC_K ,

$$\log_{10} R_K = 15.58201 - 0.4\text{BC}_K, \quad (1)$$

TABLE 4
IR PHOTOMETRY OF PROGRAM STARS

Star (1)	K (2)	$(V-K)_0$ (3)	$(J-K)_0$ (4)	BC_K (5)	M_{bol} (6)
I-012 ^a	11.33	2.65	0.69	2.23	-0.96
I-141	11.77	2.57	0.66	2.21	-0.64
I-194 ^a	11.61	3.05	0.68	2.31	-0.60
I-322	9.94	2.74	0.69	2.27	-2.31
III-152	11.52	3.03	0.71	2.35	-0.65
IV-003	11.06	2.44	0.58	2.09	-1.37
IV-025	12.46	2.37	0.52	2.00	-0.16
IV-072 ^b	11.51	3.18	0.65	2.28	-0.73
IV-167	12.58	2.81	0.60	2.17	0.23
IV-203	8.79	3.52	0.89	2.60	-3.13
IV-329	10.78	2.77	0.74	2.33	-1.41
Excess	0.14	1.40	0.26	...	...

Col. (1).—Star designation from Arp 1965.

Cols. (2)–(4).—Reddening corrections for colors and K magnitudes are entered on the last line of the table. Colors and magnitudes are from Frogel, Whitford, & Rich 1984, except for those entries with superscripts.Col. (5).— BC_K are from the correlations vs. $(V-K)$ and $(J-K)$ in Frogel & Whitford 1987.Col. (6).— M_{bol} assumes $R_0 = 8$ kpc; due to the spatial depth of the bulge, these values must be considered accurate to only ± 0.3 mag.^a Frogel & Whitford 1987.^b K and $(J-K)$ from Davidge 1991; V from Arp 1965 with Blanco & Blanco's 1984 corrections to Arp's photometry.TABLE 5
SENSITIVITY OF T_{eff} (K) TO $\Delta E(B-V) = 0.05$

T_{eff}	ΔT_{V-K}	ΔT_{V-J}	ΔT_{J-K}
4700	133	183	196
4150	99	122	155

TABLE 6
TEMPERATURES (K) FROM IR COLORS

Star	T_{J-K}	T_{V-J}	T_{V-K}	T_{AV}
I-012	4419	4505	4475	4470
I-141	4507	4564	4542	4540
I-194 ^a	4451	4105	4190	4220
I-322	4422	4402	4404	4410
III-152	4365	4162	4202	4220
IV-003	4754	4606	4649	4660
IV-025 ^a	4960	4612	4717	4740
IV-072 ^a	4540	4000	4110	4180
IV-167 ^a	4689	4233	4351	4390
IV-203	3884	3963	3959	3950
IV-329	4283	4428	4381	4370

^a Metal-rich star.

where the constant is determined from the zero point of the Johnson K -band absolute flux calibration (we used $0.665 \times 10^{-23} \text{ W m}^{-2}$, as adopted by McWilliam 1991). From the bolometric K -band corrections of FW87, and Frogel & Whitford (1982), we computed R -values for 34 warm bulge giants. Effective temperatures were obtained by interpolating the grid of R -values given by Blackwell et al. (1980). Infrared flux T_{eff} values are sensitive to reddening, but less than either $(V-K)$ or $(J-K)$ temperatures. For a K2 giant a reddening uncertainty of $\Delta E(B-V) = 0.05$ mag corresponds to an uncertainty of 95 K in T_{eff} .

FW87 quote possible BC_K errors of up to 0.07 mag, which would correspond to T_{eff} errors at the 100 K level. Assuming this level of uncertainty for the BC_K measurement and an $E(B-V)$ uncertainty of 0.05 mag, the IR flux temperatures are accurate to 138 K, on an absolute scale.

In Figures 5a and 5b we compare infrared flux temperatures with T_{eff} derived from $(J-K)$ and $(V-K)$ colors. The $(J-K)$ color temperatures agree extremely well with the IR flux method, on average 13 K hotter than the IR flux temperatures, with a rms deviation of 74 K. The $(V-K)$ color temperatures are on average 115 K below the IR flux method, with a rms deviation of 95 K. The larger rms scatter for the $(V-K)$ data is entirely due to a single, relatively large deviation for one point; without this point the rms dispersion for $(V-K)$ becomes 74

K. A 13 K uncertainty in the mean IR flux versus $(J-K)$ relation is just what would be expected from a sample of 34 points with a 1σ dispersion of 74 K. The 74 K 1σ dispersion is consistent with a 1σ dispersion in BC_K measurement of 0.05 mag, or a 5% uncertainty in total measured flux. This level of uncertainty is approximately consistent with that suggested by FW87 and with flux measurement uncertainties found for many solar neighborhood stars (Leggett et al. 1986). Using the 74 K rms dispersion of the $(J-K)$ temperatures, a 30 K uncertainty for the model atmospheres, and a 95 K uncertainty due to reddening, the absolute IR flux temperature uncertainty becomes ± 124 K. If an optimistic reddening uncertainty of 0.03 mag is utilized, this becomes ± 98 K. For relative temperatures, under the assumption of equal reddening for all stars, the uncertainty is ± 80 K.

Figure 5a indicates that the $(J-K)$ color temperature is identical to the relation found for solar neighborhood GK giants. Because the $(J-K)$ temperatures are quite sensitive to reddening (McW90), the agreement with IR flux temperatures indicates that the reddening used by FW87 is correct. We point out that there is some interdependency between the IR flux and color temperatures, such as the common reddening and the bands used to form the colors and total fluxes. We note that the 74 K rms dispersions seen in Figures 5a and 5b are consistent with $(J-K)$ and $(V-K)$ colors with rms uncertainties of 0.03 mag.

The shift between IR flux and $(V-K)$ color temperatures cannot be explained by a small reddening error, because an increased reddening adopted to increase the temperature derived from $(V-K)$ also increases the apparent IR flux temperature, with a differential rate of about 52 K for each 0.10 mag in $E(B-V)$. To explain the 115 K difference would require an increase of 0.22 mag in $E(B-V)$; however, this is completely ruled out by uncertainties in the measured $E(B-V)$ [e.g., Walker & Terndrup 1991 give $\sigma E(B-V) = 0.05$ mag].

An alternative explanation for the 115 K shift of the $(V-K)$ temperatures is a zero-point error of +0.13 mag in the V -band magnitudes; this is not unreasonable, as the only V magnitudes available are from the photographic photometry of Arp (1965). This explanation is supported by Figure 3a of Blanco & Blanco (1984), which clearly shows that for these stars, in the range $V \sim 15.0$ –16.0, the Arp (1965) V magnitudes are systematically higher than CCD measurements by about 0.1 mag, with a dispersion of about ± 0.1 mag. The larger dispersion in $(V-K)$ temperatures could reasonably be due to a glitch in the V -band measurement for the deviant star, or perhaps V -band blanketing resulting from high metallicity.

Whatever the cause for the zero-point shift of the $(V-K)$ temperatures, our comparison of color and IR flux temperatures indicates that the colors provide reliable temperatures for the bulge K giants, and that the $(J-K)$ color- T_{eff} relation, at least, is identical to that found for solar neighborhood giants. We also conclude that the uncertainty in adopted temperature for our stars is highly unlikely to exceed 150 K.

In adopting the stellar gravity parameter, we utilize the following relation:

$$\log(g/g_{\odot}) = \log(M/M_{\odot}) - \log(L/L_{\odot}) + 4 \log(T/T_{\odot}). \quad (2)$$

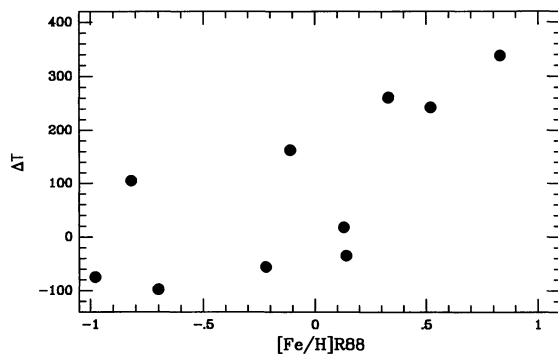


FIG. 4.—Plot of ΔT against $[\text{Fe}/\text{H}]$ from Rich (1988), where $\Delta T = T_{(J-K)} - T_{(V-K)}$. Above $[\text{Fe}/\text{H}] \sim 0.2$ dex the $(V-K)$ temperatures are, on average, 300 K cooler than temperatures derived from $(J-K)$ colors.

TABLE 7
ADOPTED ATMOSPHERE PARAMETERS

Star	T_{eff} (K)	$\log g$ (cm s^{-2}) ^a	[Fe/H] _{model}	[Fe/H] _{lines}	ξ (km s^{-1})	List
I-012	4470	1.9	-0.3	-0.34	2.0	2
I-141	4540	2.0	-0.2	-0.23	1.9	3
I-194	4450	2.0	-0.3	-0.26	1.9	2
I-322	4400	1.3	-0.2	-0.22	2.2	2
III-152	4220	1.9	-0.5	-0.49	2.0	2
IV-003	4660	1.8	-1.0	-0.94	1.7	1
IV-025	4960	2.4	0.4	0.37	1.5	3
IV-072	4540	2.0	0.0	-0.05	2.2	3
IV-167	4690	2.4	0.5	0.44	2.2	4
IV-203	3950	0.8	-1.0	-1.08	2.3	1
IV-329	4370	1.6	-0.9	-0.85	1.7	1
α Boo ^b	4280	1.3	-0.5	-0.54	1.4	1
μ Leo ^c	4540	2.3	0.5	0.42	1.8	4

^a Assuming a mass of $1.25 M_{\odot}$.
^b T_{eff} from McWilliam 1990; $\log g$ from Fe ionization equilibrium.
^c Parameters from Sneden & Gratton 1990, except for [Fe/H] and ξ .

Luminosities from Table 7 are based on the IR photometry of Frogel, Whitford, & Rich (1984), bolometric corrections from FW87, a distance modulus of 8 kpc, and reddening of $E(B - V) = 0.48$. There is some question of the masses of the stars in our sample. One could adopt a mass of $0.8 M_{\odot}$, since this is the turnoff mass of typical globular cluster systems. If we attempt to find approximate masses using the theoretical evolutionary tracks of Maeder & Meynet (1988), the metallicities of R88, and a rough dependence of the red giant branch (RGB) colors on metallicity taken from the literature, we find that the average mass of our 11 stars is somewhere near $2 M_{\odot}$. Compared with the globular cluster stars, the BW giants are either unusually blue or unusually luminous. Comparing the luminosities of our 11 K giants with typical solar neighborhood K giants, it is apparent that our sample of stars is more luminous by approximately 1 mag. This is in accord with a

higher average mass for our sample of bulge giants. However, both the metallicities and the evolutionary tracks are somewhat uncertain, and given that the RGB tracks are close together for the K giant portion of the H-R diagram, we assign a low confidence to masses determined in this manner. The apparent high luminosity of our sample of BW giants could be largely explained if the distance to the Galactic center is closer to 7 kpc than to 8 kpc (Blanco 1990) and if we systematically selected stars on the solar neighborhood side of the edge of the bulge.

Because of the uncertainty over the masses, we adopt the maximum permitted value of the bulge turnoff mass according to Terndrup (1988), at $1.25 M_{\odot}$.

Despite this difficulty, we are able to proceed, because the final abundances are fairly insensitive to the adopted mass; a factor of 2 uncertainty will not alter our results significantly.

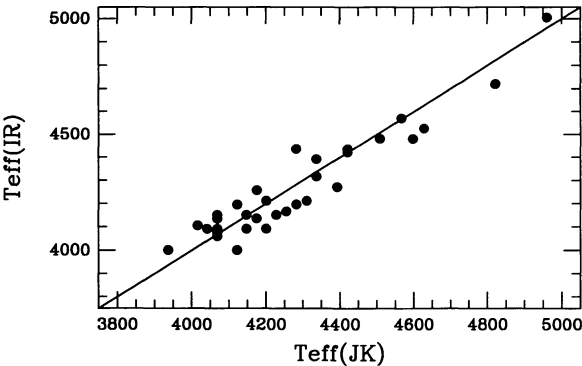


FIG. 5a

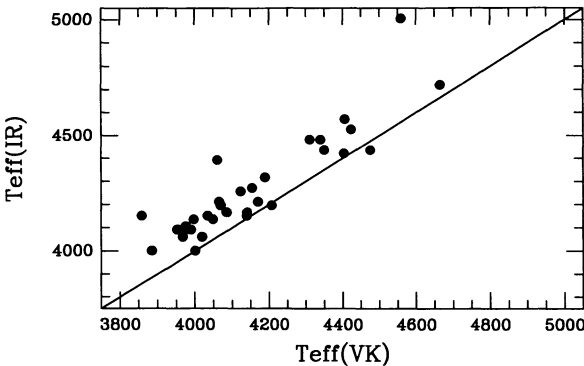


FIG. 5b

FIG. 5.—(a) Comparison of infrared flux temperatures, $T_{\text{eff}}(\text{IR})$, and effective temperatures derived from $(J - K)$ colors, $T_{\text{eff}}(JK)$. The rms dispersion of the correlation is 74 K. (b) Comparison of infrared flux temperatures, $T_{\text{eff}}(\text{IR})$, and effective temperatures derived from $(V - K)$ colors, $T_{\text{eff}}(VK)$. The rms dispersion of the correlation is 95 K, or 74 K if the most deviant point is rejected. The $(V - K)$ color temperatures are, on average, 115 K cooler than the infrared flux temperatures.

The microturbulent velocity and model atmosphere metallicity parameters were found iteratively through abundance analysis of the many Fe I lines; details are described in § 3.4.

3.3. Oscillator Strengths

Whenever possible, we adopt accurate laboratory gf values for our abundance analyses. In cases where no laboratory gf values exist, or available laboratory gf values are highly uncertain, we compute gf values based on the solar equivalent widths using the Holweger & Müller (1974) model solar atmosphere. Unfortunately, for weak lines and lines in the red portion of the spectrum the laboratory data are very sparse, and frequently one has to use multiple sources for a single element.

We have attempted to place the various sets of Fe I gf values onto a single system by intercomparison of common log gf values. First we compared the 95 lines redward of 4300 Å common to the Oxford gf values (e.g., Blackwell et al. 1982a; Blackwell, Petford, & Simmons 1982b; Blackwell, Petford, & Shallis 1979) and those given by O'Brian et al. (1991, hereafter OB91). We found that for log gf less than -2.0 the Oxford log gf values are systematically lower than the OB91 sample by 0.033 with a σ of 0.043. The relative shift of the Oxford scale is in excellent agreement with the value of 0.034 found by Bard, Kock, & Kock (1991). Because both the Bard et al. (1991) and the OB91 studies possess the same shift relative to the Oxford scale, we prefer to adjust the Oxford log gf values to the O'Brian/Bard systems. For lines with log $gf \geq -2.00$ we find no systematic shift between the Oxford and the OB91 log gf values. The scatter between the gf scales is very small, at 0.043 dex, suggesting that σ for the individual systems is of order 0.03 dex.

These gf studies are severely deficient in Fe I lines in the red region of the spectrum; therefore we supplement our log gf database with values from May, Richter, & Wichelmann (1974, hereafter May74) and Gurtovenko & Kostik (1981, hereafter GK81). The GK81 study computed gf values by matching equivalent widths of Fe I lines found in the solar spectrum, relied on the large set of laboratory gf values from May74 for the Fe abundance, and was ultimately normalized to the Oxford scale. It is of interest to note that the GK81 solar abundance derived solely from the May74 gf values was $\epsilon(\text{Fe}) = 7.525$, which is in agreement with recent determinations of the solar iron abundance (Biémont et al. 1991; Holweger, Heise, & Kock 1990; Holweger et al. 1991).

We made a comparison between GK81 and May74 gf values for 46 lines in the region covered by our spectra (6140 and 7600 Å), and found the May74 log gf values to be systematically larger than the GK81 values by 0.073 dex, with a σ of 0.050; there was no obvious slope to the plot of difference with log gf value.

Two points in the sample of lines common to the GK81 and May74 data sets stood out significantly from the general distribution, namely, Fe I at 6330.8 and 6806.8 Å: the line at 6330.8 Å can be explained by a blended solar line, so the May74 value is adopted; whereas the 6806.8 Å line cannot be explained by a blend, we suspect a typographical error for the 1 dex difference in gf values and adopt the GK81 value (this is also closer to the Kurucz 1992 log gf for the line).

With the GK81 gf 's shifted to the May74 system we com-

pared May/GK gf 's with those of lines common to the O'Brian/Bard system in the region 6140–7600 Å (which included corrected Oxford gf values). From the 40 lines in common no unambiguous shift or slope could be found between the systems; the formal error on the mean shift of 0.00 is ± 0.015 dex. This result is consistent with the results of Bard et al. (1991), who find no systematic differences with the May74 data. We find the scatter between the May74 system and the O'Brian/Bard system to be $\sigma = 0.10$ dex, in excellent agreement with Bard et al. (1991). Because of the possibility of undetected blending in the solar spectrum, we ignored any gf values from GK81 which were larger than those from other gf sources by more than 0.30 dex (3σ on the GK81 system).

Finally, weighted mean gf values were computed with the following σ values: O'Brian et al. and Bard et al. log gf 's $\sigma = 0.03$; Oxford group log gf 's $\sigma = 0.04$; May74 and GK81 log gf 's $\sigma = 0.10$.

Our log gf values for other species came from a variety of sources; where possible we adopt accurate laboratory determined values. For a number of lines there are no laboratory measurements, or only very unreliable values; we have computed log gf for a number of these lines, using equivalent widths measured from the Liège Solar Atlas (Delbouille, Roland, & Neven 1973), and the solar model atmosphere of Holweger & Müller (1974). In a few cases we take the unweighted average of the computed solar gf value and the gf value from another source. In Table 8 we list the adopted log gf values for each line and identify the literature source in each case; multiple sources indicate that we used average log gf .

A total of 37 solar gf values were computed and are identified in Table 8; 29 of these lines appear in the solar gf list of Thévenin (1990). The gf values of Thévenin (1990) were also based on EW measurement of features in the Liège Solar Atlas, but used the solar model atmosphere of Gustafsson et al. (1975). A comparison of the lines in common with Thévenin (1990) shows that our solar gf values are, on average, 0.04 dex larger, with a rms difference of 0.14 dex. Assuming an equal contribution to the rms scatter from both sets of solar gf values the rms scatter from each source is 0.10 dex. A plot of the Thévenin (1990) log gf values against those found here shows no systematic deviation from over six orders of magnitude. To test the absolute accuracy of the solar gf values, we compared the 36 Fe I lines in our sample possessing accurate O'Brian et al. (1991) laboratory oscillator strengths with the same lines in the Thévenin (1990) list. We find the O'Brian gf values to be larger by 0.06 dex; after correcting for this zero point, the rms scatter is 0.11 dex. We conclude that the Thévenin (1990) solar gf values and those computed here are accurate to a 1σ level of 0.11 dex.

CN red system wavelengths, branch identifications and quantum numbers, from the work by Davis & Phillips (1963), were kindly supplied by J. G. Phillips in floppy disk format.

To compute CN line gf values, we scaled the band oscillator strengths with the following equation:

$$f(v'v'') = \frac{R_e^2(r_{v'v''})q(v'v'')\lambda(00)}{R_e^2(r_{00})q(00)\lambda(v'v'')} f(00), \quad (3)$$

where $q(v'v'')$ is the Franck-Condon factor and $R_e(r_{v'v''})$ is the transition moment.

TABLE 8
ABUNDANCES FOR α BOO

Ion	Wavelength (Å)	EP (eV)	log <i>gf</i>	Refs.	EW (mÅ)	$\epsilon(M)$	List
Li I	6707.800	0.00	0.171	y	≤ 3	≤ -0.82	2,1
Na I	6154.230	2.10	-1.570	r	69	5.89	3,2,1
Na I	6160.753	2.10	-1.270	r	...	...	1
Mg I ^b	6319.242	5.11	-2.215	rt	52	7.37	3,2,1
Mg I	6765.450	5.75	-1.94	z			
Mg I ^a	6894.890	5.75	-2.390	t	48	8.27	4,3,2,1
Mg I	6965.408	5.75	-1.720	t	51	7.68	3,2,1
Mg I	7387.700	5.75	-0.870	t	83	7.31	4,3,2,1
Al I	6696.032	3.14	-1.481	p	92	6.51	4,3,2,1
Al I	6698.669	3.14	-1.782	p	63	6.32	3,2,1
Si I	6237.328	5.61	-1.010	t	63	7.36	1
Si I	6721.844	5.86	-1.169	t	43	7.44	3,1
Si I	6976.504	5.95	-1.044	t	32	7.21	4,3,2,1
Si I	7226.208	5.61	-1.440	t	33	7.18	4,3,2,1
Si I	7235.325	5.61	-1.460	t	27	7.07	4,3,2,1
Si I	7405.790	5.61	-0.660	t	80	7.29	3,1
Si I	7415.958	5.61	-0.730	t	78	7.35	1
Si I	7423.509	5.62	-0.580	t	98	7.58	1
Ca I	6166.440	2.52	-1.142	j	121	6.10	4,3,2,1
Ca I	6169.044	2.52	-0.797	j	138	6.10	4,3,2,1
Ca I	6169.564	2.52	-0.478	j	177	6.40	4,3,2,1
Ca I	6455.605	2.52	-1.290	j	109	5.98	3,2,1
Ca I	6508.846	2.52	-2.500	t	44	6.03	1
Ca I	6572.795	0.00	-4.310	t	171	6.36	4,3,2,1
Ca I	6719.687	2.71	-2.550	t	17	5.74	3,2,1
Ca I	7148.150	2.71	0.137	j	204	6.15	4,3,2,1
Ca I	7326.160	2.93	-0.230	t	166	6.28	4,3,2,1
Sc I	6210.671	0.00	-1.570	i	116	3.16	2,1
Sc II	6604.600	1.36	-1.480	i	87	2.95	2,1
Ti I	6303.767	1.44	-1.566	i	88	4.76	1
Ti I	6336.113	1.44	-1.743	i	82	4.84	2,1
Ti I	6508.124	1.43	-2.050	t			
Ti I	6554.238	1.44	-1.218	i	107	4.70	1
Ti I ^c	6556.077	1.46	-1.074	i	142	5.22	4,3,2,1
Ti I	6743.127	0.90	-1.630	i	143	4.90	4,3,2,1
Ti I	6861.500	2.27	-0.740	i	64	4.65	3,2,1
Ti I	7209.468	1.46	-0.500	i	173	4.97	1
Ti I	7216.190	1.44	-1.150	i	119	4.74	1
Ti I	7251.717	1.43	-0.770	i	133	4.57	4,3,2,1
Ti II	6559.576	2.05	-2.480	w	63	4.82	2,1
Ti II	7214.741	2.59	-1.740	i	48	4.49	1
V I	6274.658	0.27	-1.670	i	130	3.95	4,3,2,1
V I	6285.165	0.28	-1.510	i	122	3.63	4,3,2,1
V I	6504.186	1.18	-1.230	i	62	3.61	2,1
V I	6531.429	1.22	-0.840	i	88	3.71	4,3,2,1
V I	6543.510	1.19	-1.660	i	35	3.63	1
V I	6812.356	1.04	-2.110	i	35	3.83	1
Cr I	6330.096	0.94	-2.920	i	100	4.82	3,2,1
Cr I	6881.716	3.44	-0.450	t	60	5.11	4,3,2,1
Cr I	7400.188	2.90	-0.111	i	121	5.07	2,1
Cr I	7462.342	2.91	-0.010	t	146	5.38	1
Fe I	6151.623	2.18	-3.330	adfg	105	6.85	4,3,2,1
Fe I	6173.341	2.22	-2.863	dfg	145	7.28	2,1
Fe I	6180.209	2.73	-2.628	abfg	115	7.15	2,1
Fe I	6187.995	3.94	-1.673	fg	80	7.07	4,3,2,1
Fe I	6200.321	2.61	-2.386	acfg	124	6.90	3,2,1
Fe I	6212.067	4.37	-2.690	t	15	7.22	1
Fe I	6213.437	2.22	-2.490	ag	148	6.96	2,1
Fe I	6219.287	2.20	-2.428	adg	171	7.29	2,1
Fe I	6229.232	2.83	-2.821	bfg	93	6.98	3,2,1
Fe I	6240.653	2.22	-3.212	abfg	121	7.09	3,2,1
Fe I	6246.327	3.60	-0.796	abg	139	7.02	2,1

TABLE 8—Continued

Ion	Wavelength (Å)	EP (eV)	log <i>gf</i>	Refs	EW (mÅ)	ϵ(M)	List
Fe I ^b	6254.253	2.28	-2.435	abf	184	7.61	2,1
Fe I	6265.141	2.18	-2.532	ad	190	7.66	2,1
Fe I	6270.231	2.86	-2.543	abfg	93	6.73	4,3,2,1
Fe I	6271.283	3.32	-2.809	fg	61	6.97	3,2,1
Fe I	6297.799	2.22	-2.669	adfg	140	6.97	1
Fe I	6301.508	3.65	-0.701	bg	150	7.16	1
Fe I	6307.854	3.64	-3.270	t	18	6.95	4,3,2,1
Fe I	6311.504	2.83	-3.153	bfg	81	7.05	2,1
Fe I	6322.694	2.59	-2.438	acg	127	6.98	3,2,1
Fe I	6330.852	4.73	-1.640	f	40	7.26	3,2,1
Fe I	6335.337	2.20	-2.175	ag	187	7.26	2,1
Fe I	6336.830	3.69	-0.667	g	134	6.88	2,1
Fe I	6353.849	0.91	-6.360	t	35	6.74	1
Fe I	6355.035	2.84	-2.328	abf	161	7.86	1
Fe I	6380.750	4.19	-1.366	afg	86	7.22	2,1
Fe I	6392.538	2.28	-3.957	g	62	6.70	2,1
Fe I	6393.612	2.43	-1.505	abg	216	7.26	2,1
Fe I	6411.113	4.73	-2.210	t	14	7.16	2,1
Fe I	6411.658	3.65	-0.646	abg	156	7.18	2,1
Fe I	6419.956	4.73	-0.183	fg	98	7.02	2,1
Fe I	6421.360	2.28	-1.979	abdg	202	7.34	2,1
Fe I	6430.856	2.18	-1.954	adg	176	6.82	2,1
Fe I	6475.632	2.56	-2.929	afg	136	7.55	2,1
Fe I	6481.878	2.28	-2.985	adfg	139	7.25	2,1
Fe I	6494.994	2.40	-1.246	adg	270	7.39	1
Fe I	6498.945	0.96	-4.675	aefg	157	7.39	2,1
Fe I	6518.373	2.83	-2.397	abfg	119	7.05	3,2,1
Fe I	6533.940	4.54	-1.360	f	60	7.13	2,1
Fe I ^b	6546.252	2.75	-1.536	a	166	7.02	2,1
Fe I	6556.806	4.79	-1.720	t	23	7.02	4,3,2,1
Fe I	6569.224	4.73	-0.380	f	94	7.12	2,1
Fe I	6571.180	4.29	-2.950	t	15	7.36	1
Fe I	6574.254	0.99	-5.009	afg	114	6.90	4,3,2,1
Fe I	6575.037	2.59	-2.714	afg	119	7.01	2,1
Fe I	6591.306	4.59	-1.997	g	16	6.83	1
Fe I	6593.874	2.43	-2.377	adfg	182	7.61	2,1
Fe I	6597.571	4.77	-0.970	f	48	6.79	4,3,2,1
Fe I	6608.044	2.27	-3.939	fg	62	6.66	4,3,2,1
Fe I	6625.039	1.01	-5.277	g	92	6.77	2,1
Fe I	6627.560	4.53	-1.559	fg	44	6.98	4,3,2,1
Fe I	6646.966	2.60	-3.917	g	44	6.76	2,1
Fe I	6648.121	1.01	-5.730	t	72	6.86	2,1
Fe I	6653.911	4.14	-2.447	g	28	7.03	4,3,2,1
Fe I	6677.997	2.69	-1.395	ag	202	7.24	2,1
Fe I	6699.136	4.59	-2.117	g	16	6.95	4,3,2,1
Fe I	6703.576	2.76	-3.059	fg	80	6.80	2,1
Fe I	6704.500	4.20	-2.587	g	15	6.88	4,3,2,1
Fe I	6705.105	4.61	-1.060	t	56	6.83	3,2,1
Fe I	6710.323	1.48	-4.807	g	87	6.88	2,1
Fe I	6713.745	4.79	-1.479	fg	31	6.97	4,3,2,1
Fe I	6715.386	4.59	-1.540	f	41	6.99	1
Fe I	6716.252	4.56	-1.847	g	32	7.06	1
Fe I	6725.364	4.10	-2.227	g	38	6.97	4,3,2,1
Fe I	6726.673	4.59	-1.087	g	62	6.94	4,3,2,1
Fe I	6733.153	4.62	-1.479	fg	42	6.98	4,3,2,1
Fe I	6739.524	1.56	-4.801	bg	75	6.75	4,3,2,1
Fe I	6750.164	2.42	-2.592	abdf	146	7.13	3,2,1
Fe I	6752.716	4.64	-1.263	fg	53	7.01	2,1
Fe I	6806.856	2.73	-2.633	fg	92	6.57	3,2,1
Fe I	6810.267	4.59	-0.992	afg	66	6.93	4,3,2,1
Fe I	6820.374	4.62	-1.214	fg	69	7.25	3,2,1
Fe I ^b	6828.596	4.64	-0.843	fg	73	6.99	3,2,1
Fe I	6839.835	2.56	-3.378	fg	88	6.97	2,1
Fe I	6841.341	4.61	-0.733	fg	82	7.03	2,1
Fe I	6842.689	4.64	-1.224	fg	60	7.11	4,3,2,1
Fe I	6843.655	3.65	-0.863	fg	87	5.97	
Fe I	6851.652	1.60	-5.247	g	51	6.82	1
Fe I	6855.723	4.39	-1.747	g	41	6.93	3,2,1
Fe I	6858.155	4.59	-0.939	afg	66	6.88	4,3,2,1
Fe I	6898.307	4.22	-2.157	g	38	7.04	1
Fe I ^b	6911.522	2.42	-3.967	g	62	6.89	4,3,2,1
Fe I	6916.686	4.15	-1.359	fg	79	7.00	3,1
Fe I	6960.330	4.57	-1.907	g	25	6.98	1
Fe I	6978.862	2.48	-2.465	abdg	129	6.75	2,1
Fe I	6988.533	2.40	-3.519	fg	88	6.88	2,1

TABLE 8—*Continued*

Ion	Wavelength (Å)	EP (eV)	log <i>gf</i>	Refs.	EW (mÅ)	ϵ(M)	List
Fe I	7007.976	4.18	-1.929	fg	56	7.11	2,1
Fe I	7022.957	4.19	-1.148	fg	85	6.94	4,3,2,1
Fe I	7024.644	4.54	-1.106	afg	67	6.99	1
Fe I	7038.220	4.22	-1.214	fg	76	6.86	1
Fe I	7068.423	4.07	-1.319	fg	107	7.38	1
Fe I	7071.866	4.61	-1.627	g	40	7.06	4,3,2,1
Fe I	7072.800	4.07	-2.767	g	...	...	...
Fe I	7083.394	4.91	-1.327	g	34	7.03	3,2,1
Fe I	7090.390	4.23	-1.109	fg	87	6.97	1
Fe I	7112.170	2.99	-3.001	bfg	77	6.97	1
Fe I	7127.573	4.99	-1.177	g	36	7.03	1
Fe I	7130.925	4.30	-0.708	fg	114	7.22	1
Fe I	7132.985	4.06	-1.635	afg	65	6.82	3,2,1
Fe I	7142.517	4.93	-1.017	g	50	7.07	1
Fe I ^b	7145.312	4.61	-1.240	t	69	7.25	...
Fe I	7151.464	2.48	-3.657	g	73	6.82	1
Fe I	7155.634	4.99	-1.017	g	44	7.03	1
Fe I	7180.004	1.48	-4.707	g	101	6.96	1
Fe I	7219.680	4.06	-1.617	g	77	7.04	4,3,2,1
Fe I	7228.700	2.75	-3.307	g	77	6.93	1
Fe I	7256.142	4.93	-1.517	g	23	6.97	3,2,1
Fe I	7288.740	4.22	-1.277	g	105	7.48	1
Fe I	7396.526	4.99	-1.567	g	33	7.34	3,2,1
Fe I	7411.162	4.28	-0.287	g	119	6.83	2,1
Fe I	7418.672	4.14	-1.388	ag	68	6.73	4,3,2,1
Fe I	7420.241	5.08	-1.967	g	≤14	7.33	3,2,1
Fe I	7421.560	4.64	-1.727	g	35	7.08	4,3,2,1
Fe I	7445.758	4.26	+0.053	g	130	6.67	2,1
Fe I	7454.004	4.19	-2.337	g	34	7.08	2,1
Fe I	7461.527	2.56	-3.507	g	71	6.71	2,1
Fe I	7477.595	3.88	-2.560	t	30	6.80	4,3,2,1
Fe I	7491.652	4.28	-1.067	g	76	6.74	4,3,2,1
Fe I	7498.535	4.14	-2.177	g	42	7.00	3,2,1
Fe I	7506.030	5.06	-1.230	t	29	6.99	1
Fe I	7507.273	4.41	-1.107	g	74	6.92	1
Fe I	7531.153	4.37	-0.557	g	94	6.72	3,2,1
Fe I	7540.444	2.73	-3.777	g	48	6.81	2,1
Fe II	6416.928	3.89	-2.740	l	39	7.09	4,3,2,1
Fe II	6432.683	2.89	-3.550	k	41	6.69	4,3,2,1
Fe II	7449.338	3.89	-3.200	k	22	7.06	1
Co I	6188.998	1.71	-2.450	h	104	5.26	4,3,2,1
Co I	6678.849	1.96	-2.680	h	50	4.71	1
Co I	6814.961	1.96	-1.900	h	121	5.30	4,3,2,1
Co I	7052.870	1.96	-1.620	h	188	6.15	1
Co I	7417.390	2.04	-2.070	h	100	5.05	2,1
Ni I	6327.604	1.68	-3.150	h	109	5.94	4,3,2,1
Ni I	6339.118	4.15	-0.550	t	66	5.84	1
Ni I	6350.495	4.16	-1.810	t	17	5.98	1
Ni I ^b	6366.491	4.17	-0.950	t	120	7.42	...
Ni I	6482.809	1.93	-2.630	h	118	5.97	3,2,1
Ni I	6532.881	1.93	-3.390	h	74	5.80	1
Ni I	6586.319	1.95	-2.810	h	97	5.70	4,3,2,1
Ni I	6643.638	1.68	-2.300	h	184	6.45	4,3,2,1
Ni I	6767.784	1.83	-2.170	h	171	6.30	4,3,2,1
Ni I	6772.321	3.66	-0.980	h	72	5.69	2,1
Ni I	7110.905	1.93	-2.980	h	115	6.15	1
Ni I	7122.206	3.54	0.040	h	134	5.76	1
Ni I	7291.499	1.93	-2.780	h	153	6.59	1
Ni I ^b	7393.609	3.61	-0.270	t	117	5.79	4,3,2,1
Ni I	7414.514	1.99	-2.570	h	147	6.32	1
Ni I	7422.286	3.63	-0.140	h	120	5.74	1
Ni I	7522.778	3.66	-0.400	t	116	5.95	4,3,2,1
Ni I	7525.118	3.63	-0.520	t	82	5.34	3,2,1
Ni I	7555.607	3.85	0.110	t	114	5.67	3,2,1
Y I	6435.049	0.07	-0.820	o	34	1.16	2,1
Y II	6795.410	1.73	-1.250	u	15	1.49	4,3,2,1
Y II	7450.320	1.75	-0.780	v	17	1.07	2,1
Zr I	6762.398	0.00	-1.180	t	21	0.99	2,1

TABLE 8—*Continued*

Ion	Wavelength (Å)	EP (eV)	log <i>gf</i>	Refs.	EW (mÅ)	ϵ(M)	List
Ba II	6141.727	0.70	−0.077	q	161	1.75	4,3,2,1
Ba II	6496.908	0.60	−0.377	q	173	1.98	2,1
La II	6390.480	0.32	−1.520	u	28	0.54	2,1
La II	6774.260	0.13	−1.810	u	27	0.52	4,3,2,1
Eu II	6645.127	1.38	+0.204	m	33	0.29	4,3,2,1

NOTE.—Fe I λ6843 and Ni I λ6366 abundances differed from the mean by more than 1.0 dex, hence these lines are omitted from future abundance analyses.

^a log *gf* too low by ~0.8 dex, or line badly blended in α Boo.

^b Line is blended in some stars.

^c This line gives a suspiciously high ϵ(Ti).

REFERENCES.—(a) O'Brian et al. 1991; (b) Bard et al. 1991; (c) Blackwell et al. 1982a; (d) Blackwell et al. 1982b; (e) Blackwell et al. 1979; (f) May et al. 1974; (g) Gurtovenko & Kostik 1981; (h) Fuhr et al. 1988; (i) Martin et al. 1988; (j) Smith 1981; (k) Heise & Kock 1990; (l) Kurucz 1992; (m) Biémont et al. 1982; (n) Biémont et al. 1981; (o) Hannaford et al. 1982; (p) Buurman et al. 1986; (q) Wiese & Martin 1980; (r) Lambert & Luck 1978; (s) Wiese et al. 1969; (t) solar *gf* values from this paper; (u) Tomkin & Lambert 1986 (corrected to Anders & Grevesse 1989 abundance scale); (v) Smith 1989; (w) Spite et al. 1987; (solar *gf* values corrected to Anders & Grevesse 1989 abundance scale); (y) Andersen et al. 1984. (z) α Boo *gf* values from this paper, assuming [Mg/Fe] = 0.37.

In order to evaluate this expression, we used the theoretical Franck-Condon factors for the CN red system, computed by Dwivedi et al. (1978), which covered numerous vibrational transitions. These calculations determined *q* as a function of both vibrational and rotational quantum numbers. Dwivedi et al. (1978) also computed the rotationless *r*-centroid, *r*, for each vibrational transition considered.

The band origin wavelengths, λ(*v*'*v*"'), appearing in equation (3), were computed from the prescription and measured molecular constants of Fay, Marenin, & van Citters (1971, hereafter FMV71).

We point out that, strictly speaking, the rotational dependence of *q* should have been included in our CN *gf* calculations. However, the lines in the Davis & Phillips (1963) list all possess rotational quantum numbers less than about 60. When worst-case changes in *q*(*v*, *J*) were computed for the bands in the range 6160–7600 Å, the difference was not significant enough to warrant inclusion of the effect (although for other vibrational transitions, especially with *J* up to 100, the effect is often important).

For the CN *A*–*X* transition moments, *R_e*, we used the theoretical computations of Larsson, Siegbahn, & Ågren (1983), which tabulated *R_e* as a function of internuclear separation *r*.

Larsson et al. (1983) also computed band oscillator strengths for 19 bands of the CN red system. Using their 0–0 band oscillator strength of $f_{00} = 3.25 \times 10^{-3}$ we were able to confirm, to the accuracy of the published numbers, the band oscillator strengths of all other bands, using equation (3), and the Franck-Condon calculations of Dwivedi et al. (1978). The ability to correctly predict the Larsson et al. (1983) band oscillator strengths suggests that that we have properly calculated the band oscillator strengths of vibrational transitions not covered by Larsson et al. (1983).

Line *gf* values were computed from the band oscillator

strengths by the following:

$$gf = \frac{1}{2} \mathcal{S}_{J',J''} \frac{\lambda_{v',v''}}{\lambda_{\text{line}}} f(v'v''), \quad (4)$$

where $\mathcal{S}_{J',J''}$ is the Hönl-London factor for the transition, computed using a program written by Whiting (1973).

In addition to line *gf* values for CN, it was necessary to compute wavelengths of the ¹³CN lines and excitation potentials for the lower level of each line. Because it is always best to rely upon observed, rather than computed, wavelengths, we used the molecular constants of FMV71 to compute the ¹³CN isotopic shift for each line; these were then applied to the observed (Davis & Phillip 1963) ¹²CN wavelength. We compared our ¹³CN wavelengths with the laboratory wavelengths of Wyller (1966), for several branches of the 2–0 and 3–1 vibrational transitions, with lines falling in the region 7900–8800 Å, and found a rms deviation of only 0.05 Å.

We find a rms deviation of 0.16 cm^{−1} from a sample of 20 ¹²CN *PQR* branch lines in the 2–0 band, at 7873 Å; this is somewhat larger than claimed by FMV71, who found that residuals were always less than 0.10 cm^{−1}. We are at a loss to understand this difference. We find that residuals increase with rotational quantum number (possibly with the square of *J*''). There appears to be no dependence of the residual with the vibrational quantum number. The *Q*-branch lines are most severely affected, with an average zero-point shift of 0.39 cm^{−1}. These observations of the residuals and our extensive efforts at checking with a hand calculator suggest that there is a small problem in the published constants, perhaps due to round-off errors made by FMV71. Anyway, although the residuals are too large to be used for synthesis of high *J*'' lines, the overall accuracy is good enough to compute excitation potentials and isotope shifts.

The excitation potentials were computed from the molecular constants of FMV71 using their formulae to compute energy levels.

3.4. Abundance Analysis

Abundances derived from equivalent widths and profile fitting were computed using a modified version of the spectrum synthesis program MOOG (Snedden 1973). Model atmospheres of solar metallicity or less were built by linear interpolation of the grid of Bell et al. (1976) (35 levels per model); interpolation was performed in (T_{eff} , $\log g$, $[\text{Fe}/\text{H}]$) parameter space. For the metal-rich stars appropriate model atmospheres were linearly interpolated from a metal-rich grid (including models at $[\text{Fe}/\text{H}] = 0.25$ and 0.50), kindly provided by Gustafsson (1990). Whenever possible we computed abundances using line equivalent widths rather than profile matching, because profile matching is less quantitative, is slower, and requires human judgment for the best fit.

Peterson (1992) suggested that metallicities computed from line EWs can give severely erroneous abundances in the bulge K giants, due to blending with CN lines. Given the R88 bulge star metallicities, of up to 7 times the solar value, it seemed quite likely that even normally unimportant molecules could contribute significant opacity. This would be especially true for the nonhydrides, where the column density of absorbers increases with the square of the metallicity. Therefore, we set about to determine which of our measured lines were seriously blended with CN lines.

We performed a spectrum synthesis of the entire 1460 Å with a wavelength step of 0.01 Å, at each step including opacity from all lines within 0.4 Å. The synthetic spectrum was then convolved with a Gaussian slit profile function of $R = 17,000$. Synthetic spectra were computed for μ Leo, BW IV-025, and a model with $[\text{M}/\text{H}] = -0.3$ and $T_{\text{eff}} = 4540$ K. For each model two syntheses were performed, one including atomic lines measured here plus CN lines, and one synthesis with atomic lines only. Atomic lines that were not measured in this work were not included in the syntheses. In Figures 6a and 6b we present the spectrum syntheses, including lines of ^{12}CN and

^{13}CN only, for μ Leo and BW I-194; the figures demonstrate that CN does indeed contribute significant opacity, even at $[\text{Fe}/\text{H}] \sim -0.2$ and that CN blending with the atomic lines must therefore be considered.

In order to identify contaminated atomic lines we simply overplotted the syntheses with and without CN, and visually inspected every atomic line. We did not require that every atomic line used be completely clear of all CN blends; rather, we accepted a line so long as the effect of CN would not alter the abundance significantly. Usually the acceptable blending feature strength amounted to less than 10% of the unblended line EW. The final column of Table 8 indicates the quality of each line: list 4 includes only those lines which are clean in the most severely line blanketed star (μ Leo); list 3 gives clean lines in BW IV-025; list 2 identifies lines clean at $[\text{M}/\text{H}] = -0.3$ and $T_{\text{eff}} = 4540$ K, while list 1 is the full list of lines less those lines which show systematic problems. The clean-line lists of Table 8 should be of value for those wishing to perform abundance analysis of metal-rich K giant stars. In Figure 7 we illustrate the dramatic effect of the CN blending on the derived iron abundance for μ Leo.

In order to find the best value of the microturbulent velocity parameter, we demanded agreement between the iron abundance derived from unsaturated (weak) and saturated (strong) lines. It should be noted that the very strongest lines are sensitive to conditions high in the stellar atmosphere, near to the boundary. In the abundance calculations, lines with EWs much more than 150 mÅ have significant opacity in the very first level of the model atmosphere and so are sensitive to the numerical treatment of the first layer. Therefore, the microturbulent velocity and model metallicity were adopted using Fe I lines weaker than about 150 mÅ, with emphasis on lines of 120 mÅ or less. Abundances were computed iteratively, varying the microturbulent velocity parameter, until a plot of Fe abundance against EW showed no correlation of EW with abundance. Once the microturbulence had been determined, the metallicity of the model was compared with the metallicity derived from the Fe I lines; if the model and lines differed by more than 0.1 dex, a new model atmosphere was computed

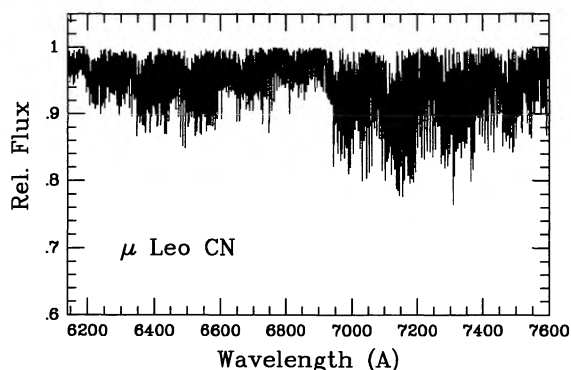


FIG. 6a

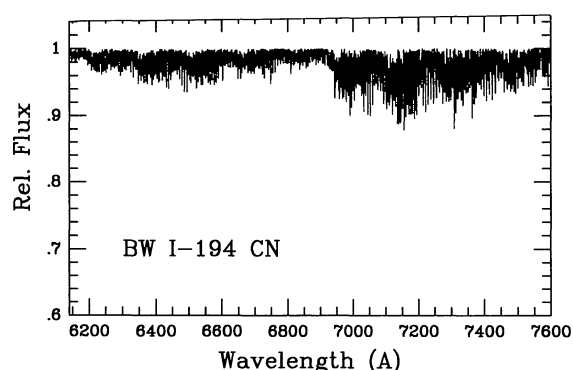


FIG. 6b

FIG. 6.—(a) Synthetic CN spectrum for μ Leo from 6140 to 7600 Å; step size was 0.01 Å. Over 8000 ^{12}CN and ^{13}CN lines were included in the calculations; lines from other molecules and atomic transitions were excluded. This demonstrates the need for careful selection of lines used in equivalent width abundance analysis. (b) Same as (a), but for BW I-194, with $[\text{Fe}/\text{H}] = -0.26$. Even below solar metallicity CN features can affect abundance results based on line equivalent widths.

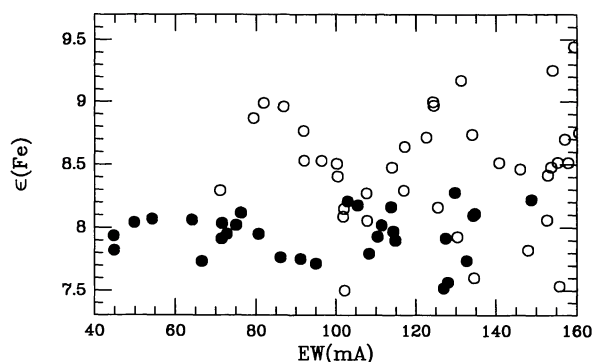


FIG. 7.—Plot of Fe abundance against equivalent width for Fe I lines in μ Leo. Open circles indicate all lines with measured EWs; filled circles indicate lines deemed to be unaffected by CN blending. If blended lines were used, the mean iron abundance would increase by ~ 0.3 dex, and the rms dispersion would be increased greatly.

using the metallicity from the Fe I lines. Iteration continued until microturbulence, model metallicity, and Fe I abundance were consistent.

The adopted T_{eff} values were checked for consistency with the Fe I abundances. For correct model atmosphere temperatures the computed Fe I abundances should be independent of

line excitation potential (EP). At the relatively low S/N and resolution of our data this method has an accuracy of only ± 400 – 500 K. Figures 8a–8d show plots of iron abundance against EP for two bulge stars and the two standards. For the standard stars, α Boo and μ Leo, it is seen that the low-EP lines give a lower abundance than the lines near 5 eV, the difference being approximately 0.10 dex. This observation is in agreement with the effects of non-LTE on Fe I lines in GK giants found by Steenbock (1985). We note that Figure 8d suggests that the adopted temperature for BW IV-167 should be lowered somewhat, which would result in a lower derived abundance. However, given the noise and low accuracy of the EP method, we choose to retain the temperatures based on IR photometry.

The model atmosphere gravities were also checked for consistency with the Fe ionization equilibrium; the iron abundances derived from Fe I and Fe II lines should be equal if the correct gravity and temperature have been adopted. The iron abundances from Fe II lines found using the spectrum profile matching method are discussed in § 3.7.

3.5. Line Profile Matching

Because of CN blanketing, especially in the SMR stars, some lines of great interest were blended, and abundances can be found only by spectrum synthesis and profile matching.

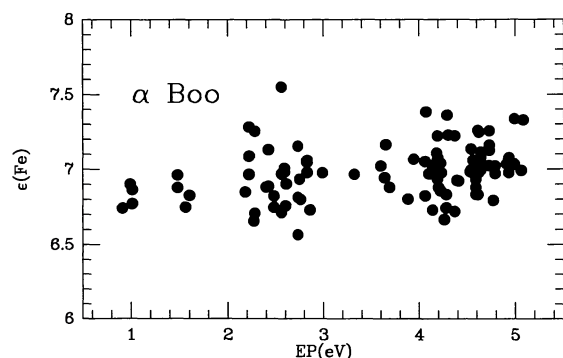


FIG. 8a

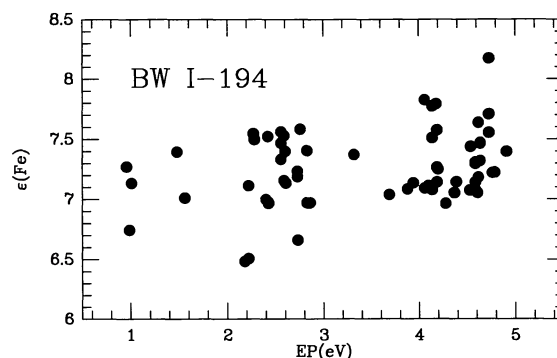


FIG. 8b

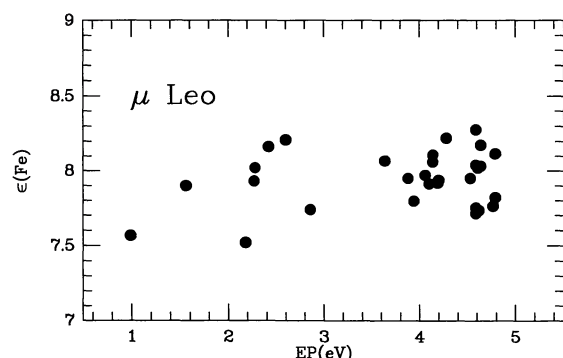


FIG. 8c

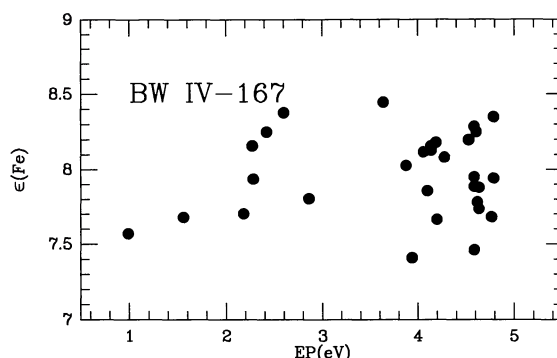


FIG. 8d

FIG. 8.—(a) Plot of iron abundance against excitation potential for Fe I lines in α Boo. In LTE the correct temperature is indicated by a zero slope in this plot. Non-LTE calculations of Steenbock (1985) show that a small decrease (0.05–0.10 dex) in abundance derived from low excitation potential lines is expected. (b) Same as (a), but for BW I-194. A small upward slope is indicated, but this is not statistically significant. (c) Same as (a), but for μ Leo. (d) Same as (a), but for BW IV-167. A small downward slope is suggested (resulting in a lower T_{eff}) but is not statistically significant.

Because of the importance of the $[\alpha/\text{Fe}]$ ratio as a diagnostic for Galactic chemical evolution, we performed spectrum synthesis for a few lines of Mg, Si, Ca, Ti, and O in all stars where telluric absorption did not affect the features. We included both ^{13}CN and ^{12}CN in our syntheses. The ^{13}CN wavelength shifts from the ^{12}CN lines were computed according to the constants of FMV71, and then applied to the observed list of CN wavelengths in Davis & Phillips (1963). The ^{13}CN line gf values used were the ^{12}CN line gf values scaled down by a factor of 20, which was the observed $^{12}\text{C}/^{13}\text{C}$ ratio found for μ Leo by Sneden & Gratton (1990). In reality the ^{13}CN line gf values are only approximately equal to the ^{12}CN line gf values; however, because the typical strength of ^{13}CN lines in μ Leo is only 10 mÅ, we can safely assume that this approximation is lost in the noise of assumptions and uncertainties.

Figures 9–13 show syntheses for a variety of α -element lines in a few of our BW stars.

3.5.1. Magnesium

Magnesium abundances are particularly difficult to determine; the most reliable indicator in our wavelength region is the 6319 Å triplet. This triplet contains the only lines covered

by our spectra for which Lambert & Luck (1978) determined the solar Mg abundance. Accurate laboratory or theoretical gf values for Mg I lines are far from complete; an excellent source is the theoretical gf work of Froese-Fischer (1975a, b). Solar gf values are difficult to compute for many of the high-excitation Mg lines, because these lines are often so strongly broadened that it is impossible to measure an accurate solar equivalent width.

We note that, although it is a problem in the solar spectrum, the effect of damping is not serious in our low-gravity program stars, owing to the density dependence of the line-broadening forces.

Initially, we attempted to find solar gf values for Mg lines present in the Sun; solar gf values were found for lines at 7387.7, 6965.4, 6894.9, and 6319.2 Å.

$\lambda 6319$.—The two lines in the 6319 Å triplet, at 6319.2 and 6318.7 Å, give solar gf values in excellent agreement with the Froese-Fischer (1975a, b) determinations. The solar Mg abundance derived for 6318.7 Å is 0.05 dex higher than the mean for BW IV-025, which we suspect is due to blending with a weak Fe II line. However, these two lines do not give concordant results for the Arcturus Mg abundance. Inspection of the

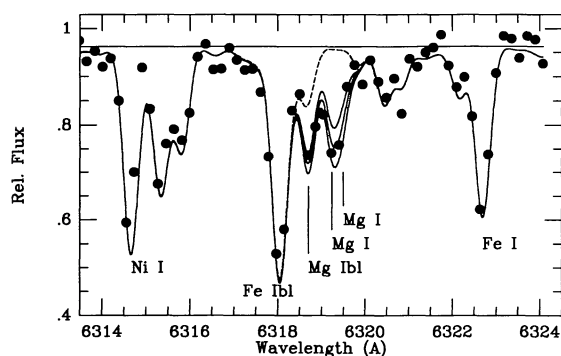


FIG. 9a

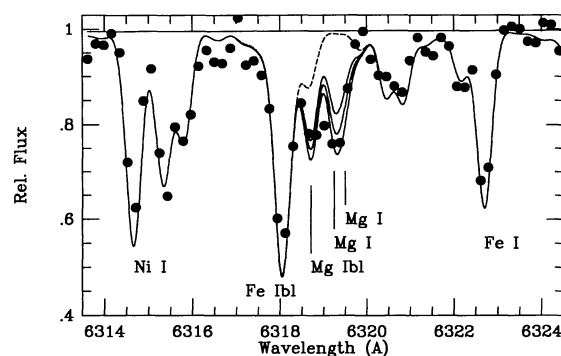


FIG. 9b

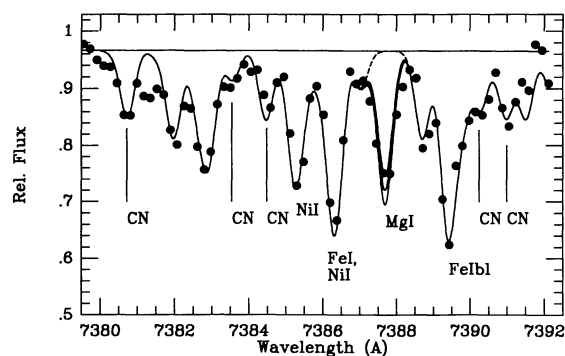


FIG. 9c

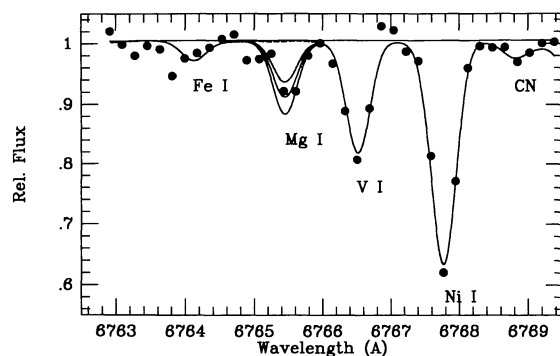


FIG. 9d

FIG. 9.—(a) Spectrum synthesis about the 6319 Å Mg I triplet region in BW I-141. Dashed line indicates a synthesis with no Mg; solid lines indicate syntheses with $[\text{Mg}/\text{Fe}] = 0.2, 0.4$, and 0.6 dex. Solid horizontal line indicates local continuum level. (b) Same as (a), but for BW I-194. (c) Spectrum synthesis about the 7387 Å Mg I region in BW IV-025. Dashed line indicates a synthesis with no Mg; solid lines indicate synthesis with $[\text{Mg}/\text{Fe}] = -0.2, 0.0$, and 0.2 dex. Solid horizontal line indicates local continuum level. Notice the good fit to the CN features in the region. (d) Spectrum synthesis about the 6765 Å Mg I line in BW I-194. Dashed line (coincident with the continuum) indicates a synthesis with no Mg; solid lines indicate syntheses with $[\text{Mg}/\text{Fe}] = 0.2, 0.4$, and 0.6 dex. Solid horizontal line indicates local continuum level. The Mg line is free from known blends. The good fit to the V I, Ni I, and CN features shows that mean abundances from Table 11 are correct.

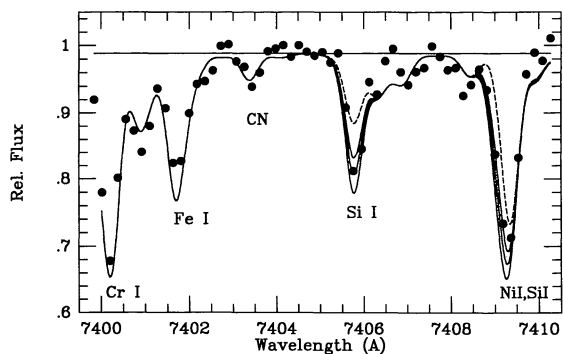


FIG. 10a

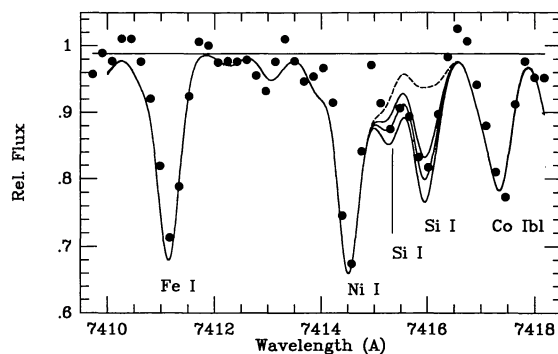


FIG. 10b

FIG. 10.—(a) Spectrum synthesis of Si I 7405 Å region in BW I-194. Dashed line indicates Si absent; solid lines are for $[\text{Si}/\text{Fe}] = -0.3, 0.0, \text{ and } 0.3$. Notice the good fit to the CN feature. (b) Spectrum synthesis of Si I 7415 Å region in BW I-194. Dashed line indicates Si absent; solid lines are for $[\text{Si}/\text{Fe}] = -0.3, 0.0, \text{ and } 0.3$.

Kurucz (1992) line list shows an Fe II line blended with the line at 6318.7 Å; increasing the Kurucz $\log gf$ value of this Fe II line from -6.06 to -4.00 forced agreement of the abundances deduced for the pair of lines in the Arcturus spectrum. That the line at 6318.7 Å gives a good Mg abundance for the Sun but a very poor abundance in giants (like Arcturus and the program stars) is consistent with the behavior of an ionized line. The large relative change in gf value for the Fe II contaminant is reasonable, because theoretical $\log gf$ values for very weak lines are highly sensitive to small changes in the wave functions. However, because of the blending we do not recommend use of the 6318.7 Å member of the triplet to determine Mg abundances.

At the resolution of the program star spectra the lines at 6319.2 and 6319.5 Å are blended together, so spectrum synthesis is necessary for abundance analysis. In Figures 9a and 9b we present spectrum synthesis fits for the 6319 Å region in BW I-141 and BW I-194.

$\lambda 6965$.—This line is blended at the resolution of our spectra, and spectrum synthesis was therefore necessary. No laboratory oscillator strengths are available for this line, so a value based on the strength of the line in the Liège Solar Atlas is used.

However, this line gives unusually large Mg abundances, inconsistent with remaining Mg I lines. Until the source of the discordant behavior of this line can be accounted for, it cannot be used as a Mg abundance indicator.

$\lambda 6894.9$.—This line is relatively weak, embedded within the blanket of strong telluric absorption lines and severely blended in the red wing with a line of comparable strength. Because of these problems this line is not recommended for abundance analysis.

$\lambda 7387$.—This line is clear of CN blends, and no significant atomic contaminants interfere with the core of the line. However, the Liège Solar Atlas shows that there are weak features in the blue and red wings of this line, which are not apparent in the Arcturus atlas (Griffin 1968). These contaminating features affect the measured EW for this line in high-metallicity stars by increasing the FWHM of the line. The effect is most pronounced for metal-rich stars when the Mg I line is saturated. Because of this difficulty, spectrum synthesis of this line is appropriate for higher metallicity stars. In the higher metallicity stars we note that the strength leads to a greater sensitivity to microturbulent velocity, and so the Mg abundance derived from this line should receive lower weight in such cases. Figure

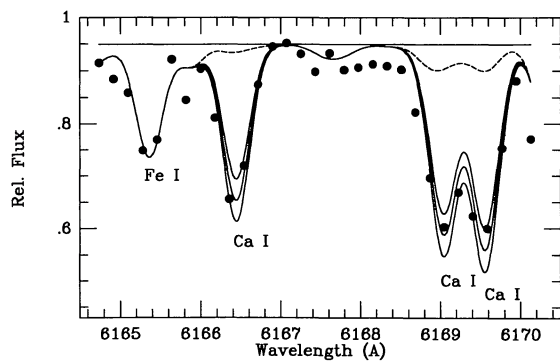


FIG. 11a

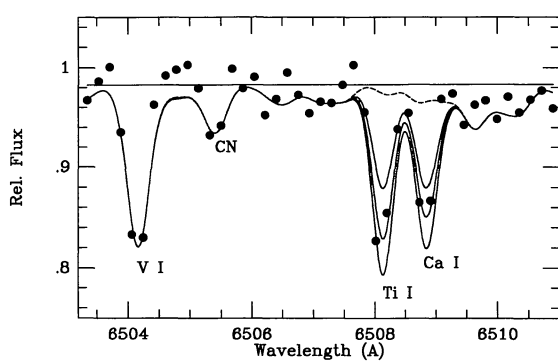


FIG. 11b

FIG. 11.—(a) Spectrum synthesis of Ca I at 6166 Å in BW I-141. Dashed line indicates no Ca; solid lines are for $[\text{Ca}/\text{Fe}] = -0.3, 0.0, \text{ and } 0.3$. Weak CN features contaminate all three Ca lines in this region. (b) Spectrum synthesis of Ca I at 6508 Å in BW I-141. Dashed line indicates Ca and Ti absent; solid lines are for $[\text{Ca}/\text{Fe}] = 0.0, 0.2, \text{ and } 0.4$ and $[\text{Ti}/\text{Fe}] = 0.0, 0.3, 0.5$. Notice the good fit to the CN feature.

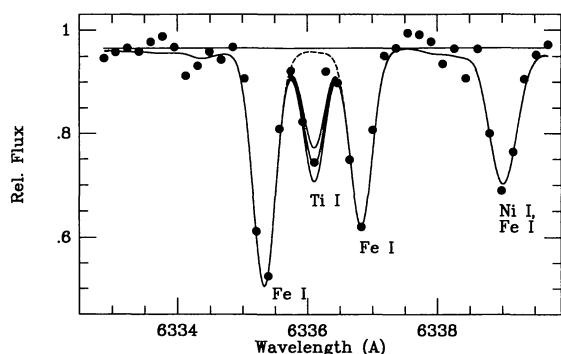


FIG. 12.—Spectrum synthesis for Ti line at 6336 Å in BW I-141. Dashed line represents no Ti; solid lines indicate $[Ti/Fe] = 0.3, 0.5$ and 0.7 .

9c shows this line and the spectrum synthesis in BW IV-025; notice the good match to the CN features in the synthesis. The 7387 Å Mg I line, due to its strength, should be an excellent Mg indicator in metal-poor giants.

Because of the uncertainty in derived abundances, we attempted to increase the number of Mg lines by using gf values based on the Arcturus spectrum (Griffin 1968). In order to find Arcturus gf values, we first needed to find the Arcturus Mg abundance; from the weak Mg lines studied by Mäcke et al. (1975) we concluded that $[Mg/Fe] = 0.37$. This value agrees well with the result derived from our lines at 6319.2, 7387.7, and 7657.6 Å; it also is in excellent agreement with the $[Mg/Fe]$ ratio of 0.4 found by Peterson, Dalle Ore, & Kurucz (1993). Using EWs measured from the Griffin atlas, we found gf values for the lines; however, we found only one usable line, at 6765.4 Å, with an Arcturus $\log gf$ value of -1.94 .

In Figure 9d we present the 6765 Å region of BW I-194, with spectrum syntheses for $[Mg/Fe] = 0.20, 0.40$, and 0.60 dex.

3.5.2. Silicon

Oscillator strengths for the Si I lines were computed from EW measurements of the lines appearing in the Liège Solar Atlas (Delbouille et al. 1973). The best laboratory measure-

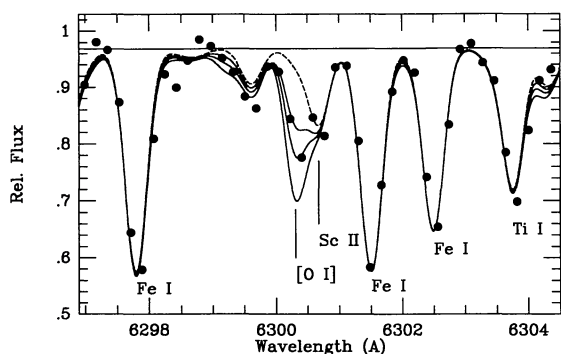


FIG. 13.—Spectrum synthesis for [O I] region at 6300 Å in BW I-012. Dashed line indicates no O, C, or N; solid lines are for $[O/Fe] = 0.0, 0.3, 0.6$. The line in the red portion of the feature is due to Sc II. For all solid line syntheses $[C/Fe]$ and $[N/Fe]$ were set to 0.0, the solar ratio. The effects of different O abundances on CN features, via the CNO equilibrium, are clearly visible at 6299.7 and 6304.1 Å.

ments of Si I $\log gf$ values are those of Garz (1973), but they had to be revised upward by 0.10 dex (see Becker, Zimmermann, & Holweger 1980). Although the revised Garz (1973) gf values fix the solar silicon abundance at 7.55 ± 0.05 , the uncertainties on individual Si I $\log gf$ values for lines in the red range from ± 0.08 to ± 0.13 dex. Given the necessity of applying offsets to the Garz $\log gf$ values, and uncertainties equal to or slightly larger than can be obtained using solar $\log gf$ values, we decided to adopt the solar $\log gf$ values.

$\lambda\lambda 6237.3$ and 6244.5 .—These lines are useful as Si abundance indicators for stars with $[Fe/H]$ less than about -0.4 dex. In the same region exists a line at 6243.8 Å, but this appears to be blended with an unidentified line and the wing of a strong V I line.

$\lambda 6721.8$.—This line is the strongest member of a triplet, the other lines being at 6719.5 and 6720.8 Å. Unfortunately, the line at 6719.5 Å is severely blended in the Sun with an unidentified line of equal strength and is not a reliable abundance indicator in our stars. The weak feature at 6720.8 Å is blended and is very insensitive to the Si abundance. Sneden & Gratton (1990) claim that the 6721.8 Å line is a reliable Si abundance indicator. However, we note that this line gives suspiciously high Si abundances in metal-rich stars, relative to the Si abundance derived from other lines.

$\lambda 6976.5$.—Although used by Lambert & Luck (1978) for the solar Si abundance, at our resolution this line is blended with an Fe I line at 6976.2 Å; furthermore, the wavelength of the feature, as measured in all of our spectra, is coincident with the Fe I line. We conclude that this line is too blended to be of any use in our spectra.

$\lambda 7226.2$.—This Si line is the only useful member of a triplet. In the solar spectrum there are two lines blueward of the Si I line near 7226.0 Å, which should be included in the synthesis. The solar lines are coincident with Ti I $\lambda 7226.02$ and Fe I $\lambda 7225.83$, as identified in the Kurucz (1992) atomic line list; however, the solar gf value found for the Ti I line is too large, and results in an overwhelmingly strong feature in a BW IV-025 synthesis. We conclude that the solar line is not due to a pure Ti I line, and so we omit the solar information in our synthesis of this region. Because of the unidentified blend, the measured wavelength of the Si region in our SMR stars is always bluer than the true wavelength of the Si I line.

$\lambda 7235.3$.—Unfortunately, this line is a thicket of telluric absorption lines, and for all of the metal-rich stars, including μ Leo, it is obliterated by the telluric.

$\lambda 7405.8$.—This region includes two Si I lines, at 7405.8 and 7409.6 Å. Both lines are significantly blended, the 7405 Å line with CN and the 7409 Å line with a Ni I line at 7409.4 Å. These blends make the features less sensitive to the Si abundance, and the resulting Si abundances more sensitive to errors in continuum placement, than other Si I lines. Figure 10a shows synthesis for this region with no Si (dashed line) and with $[Si/Fe] = -0.3, 0.0$, and 0.3 dex. One cannot appeal to photon noise to increase the abundance for the 7409 Å line, but continuum placement, Ni abundance, or Ni I gf value could affect the best-fit abundance.

$\lambda 7034$.—This line is relatively free from blends and so should be a good indicator for the Si abundance, although synthesis is still required for the SMR stars. We are unable to identify the source of a feature blueward of the Si I line; how-

ever, the unidentified feature should not affect the core of the Si I profile in any way.

$\lambda 7415$.—Synthesis is again necessary for this double Si I line feature in metal-rich stars. The line at $7415.9 \text{ \AA}$ is blended with CN and a Cr II line; only the Kurucz gf value is available for this Cr II line. The data and synthetic spectrum for this line in BW I-194 are shown in Figure 10*b*.

$\lambda 7423$.—This line is severely affected by strong CN absorption and so can only be used as a Si abundance indicator in relatively metal-poor stars, with $[\text{Fe}/\text{H}]$ less than about -0.4 dex.

3.5.3. Calcium

$\lambda 6166.4$.—This line is the cleanest of several Ca I lines in the region. Unfortunately, it is relatively strong, resulting in a high sensitivity to the microturbulent velocity parameter. See Figure 11*a* for a synthesis of this region in BW I-141.

$\lambda 6464.7$.—We believe this line to be misidentified in the Sun. A solar gf value, obtained from the Liège EW of the line, yields a significant deficiency of the line in synthesis of the BW IV-025 spectrum. However, when the Kurucz (1992) line list is included, a good match is obtained to the observed spectrum. It seems that the feature is, in fact, due to Fe I. This conclusion is bolstered by the good agreement between solar and Kurucz gf values for the feature, assuming it is all due to Fe I, and the improved fit to the BW IV-025 spectrum obtained when the solar gf value is used. We therefore eliminate this line from the list of useful Ca abundance indicators.

$\lambda 6455.6$.—Relatively clear from CN, although synthesis is necessary. Continuum placement is particularly good in this region.

$\lambda 6417.7$.—Syntheses with the Kurucz (1992) line list indicate that this feature is entirely due to the Co I line at $6417.78 \text{ \AA}$; however, the solar wavelength measured from the Liège atlas is in better agreement with the Ca I line at $6417.70 \text{ \AA}$. If synthesis calculations are performed excluding the Co I line, this feature indicates $[\text{Ca}/\text{Fe}] = 0.8$ in BW IV-025. In α Boo the Ca abundance derived from this line is ~ 0.7 dex higher than the mean from reliable lines. We conclude that until the oscillator strengths for all lines in this region are accurately determined, the feature is an unreliable Ca abundance indicator.

$\lambda 6508.9$.—The region is blanketed with many weak CN lines, but the feature is a reasonably good Ca abundance indicator. A Ti I line nearby is useful for the Ti abundance. A synthesis of this region in BW I-141 is shown in Figure 11*b*.

$\lambda 6709.9$.—This line was used by Sneden & Gratton (1990) for abundance analysis of μ Leo. Unfortunately, at our relatively low resolution this line is blended with a strong Fe I line at $6710.3 \text{ \AA}$ and is a relatively weak feature.

It is a pity that medium-strength Ca I lines are so sparse in our region; future investigations might pursue lines at bluer wavelengths.

3.5.4. Titanium

Titanium is well represented by many medium-strength lines from the neutral species; two or three Ti II lines, of uncertain quality, are also available in our spectra.

$\lambda 6336.1$.—This is a weak Ti I line in stars near solar metallicity, and is flanked by Fe I lines in both red and blue wings, which necessitates spectrum synthesis, as shown in Figure 12.

$\lambda 7496.1$.—This line is apparently clear from all blends, and is of medium strength in stars near solar metallicity. These qualities, and the plentiful local continuum, make this an important Ti abundance indicator.

3.5.5. Oxygen

Only the metastable [O I] line at $6300.3 \text{ \AA}$ is of potential use for abundance analysis in our spectra. At our relatively low resolution the line is blended with a Sc II line at $6300.7 \text{ \AA}$, and synthesis calculations are therefore required. Analysis of the region is complicated by the telluric [O I] emission line. When the stellar radial velocity is low, we abandon attempts for analysis of the stellar oxygen feature because reliable sky subtraction of the extremely strong telluric [O I] line is not possible to the level of accuracy required. Another difficulty is presented by a small number of fairly weak ($\sim 20 \text{ m\AA}$) telluric absorption lines in the region. Since we did not take telluric standard spectra before and after each stellar spectrum, we prefer to avoid analysis of the [O I]/Sc II blend when the stellar radial velocity brings this feature into coincidence with one of the telluric absorption lines.

In Figure 13 we present syntheses for the [O I] region in BW I-012.

3.5.6. Europium

In our spectra there are only two europium lines of potential use, at 6437.7 and $6645.1 \text{ \AA}$. Future work employing higher resolution spectra at high S/N with larger spectral coverage might increase the number of useful Eu lines.

Spectrum synthesis of these lines is necessary to find the most precise abundance results, because the lines are split into hyperfine components, resulting in a larger EW than would be expected from a single line alone. However, spectrum synthesis is only strictly necessary when the lines are strong enough that saturation effects become important.

$\lambda 6437$.—Synthesis of this region indicates that the Eu II line at $6437 \text{ \AA}$ is so severely blended that it is of no practical use for Eu abundance determinations in our spectra.

$\lambda 6645$.—This line is close to a weak Fe I line in the solar spectrum; at the resolution of our spectra this Fe I line is blended with the red wing of the Eu II line, and spectrum synthesis is therefore necessary for abundance analysis. There is also a weak CN feature which can slightly affect the red wing of the europium line if the star has significant CN opacity. In our synthesis the hyperfine components listed by Steffen (1985) are used; although perfectly adequate for a first-order hyperfine structure (hfs) treatment, the Steffen (1985) hfs components are only approximate and should not be used for abundance analysis of very strong lines.

Using high-resolution spectra (e.g., McW90) it is possible to measure the Eu line alone, obviating the need for spectrum synthesis when the line is unsaturated. Figure 14 shows this line in BW IV-072.

3.6. Ionization Equilibrium

A spectroscopic check on the gravity can be obtained by comparing abundances derived from ionized and neutral lines, such as Fe I and Fe II, Ti I and Ti II. The iron abundance derived from Fe II lines is also important for comparison with

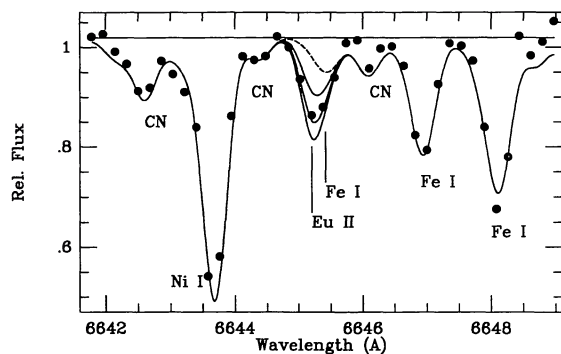


FIG. 14.—Spectrum synthesis for the Eu II region at 6645 Å in BW IV-072. Dashed line indicates no Eu present; solid lines are for synthesis at $[\text{Eu}/\text{Fe}] = 0.0, 0.4, \text{ and } 0.6$. The effect of the nearby Fe I blend is clear.

elements which are only represented by ionized lines, such as the heavy metals Eu, Ba, and Y. The ratio of these elements to iron is of great significance to chemical evolution theory. For these important abundance ratios it is desirable to remove any possible systematic effects resulting from errors or uncertainties in the model atmosphere; this can be performed by ratioing abundances derived from these ionized heavy-element lines with the iron abundance derived from Fe II lines. Unfortunately, very few good Fe II and Ti II lines can be found in our spectral window, especially for the most metal-rich stars.

Fe II is represented by five potentially useful lines, at 6247.6, 6416.9, 6432.7, 6516.1, and 7449.3 Å. In synthesis calculations with a model for our SMR bulge giant BW IV-025, we find that only the lines at 6247.6 and 6432.7 Å are clear from blends. The only difficulty with the 6432.7 Å line is the difference of 0.14 dex between the laboratory (Heise & Kock 1990) and solar gf values. We find a solar $\log gf$ value of -2.40 for the 6247.6 Å line. The lines at 7449.3 and 6516.1 Å are severely blended with CN features and only useful in stars with $[\text{Fe}/\text{H}]$ less than about -0.5 dex. The reliability of the line at 6416.9 Å is uncertain, because the Kurucz (1992) line list indicates that this Fe II line is blended with a line of Fe I with almost the same wavelength. The synthetic profile for the feature in BW IV-025 is much too deep, indicating an error in the gf values for at least one of the two lines. In fact, removing either one of the two lines does not weaken the theoretical profile enough to bring agreement between the data and synthesis for BW IV-025. There is good agreement between laboratory gf and the solar gf value, assuming that the solar line is entirely due to Fe II, which could be used to justify elimination of the Kurucz Fe I line; however, the Kurucz Fe I line is quoted with such a large gf value (at ~ 0.1) that it is difficult to understand how Kurucz could have made such an error of one or two orders of magnitude. Given this uncertainty, it is presently inappropriate to use the 6416.9 Å line as an abundance indicator for Fe II.

We adopt Fe II abundances from spectrum synthesis of the regions about the lines at 6432.7 and 6247.6 Å.

Inspection of Tables 11A–11M indicates that in most cases the Fe I and Fe II abundances differ by less than 0.2 dex, suggesting that the adopted gravities are approximately correct. However, for BW IV-025 and BW I-322 $\Delta\epsilon(\text{Fe}) = \epsilon(\text{Fe I}) - \epsilon(\text{Fe II})$ are 0.44 and 0.29 dex, respectively, and suggest that we

have adopted low $\log g$ values for these stars. The average and standard deviation of $\Delta\epsilon(\text{Fe})$ is -0.03 and 0.13 dex, respectively, for all stars excluding BW IV-025 and BW I-322, and 0.04 and 0.20 dex, respectively, for all 11 bulge stars.

Abundances from the Ti II lines at 6559.6 and 7214.7 Å enable a check on the ionization equilibrium for eight of the 11 BW giants. The mean and standard deviation of $\Delta(\text{Ti}) = \epsilon(\text{Ti I}) - \epsilon(\text{Ti II})$ are -0.10 and 0.16 dex, respectively; only BW IV-203 has a large $\Delta(\text{Ti})$, at -0.32 dex. As with the Fe ionization equilibrium, there seems to be a tendency toward larger ionized abundances than neutral line abundances. This might be interpreted as being due to a small systematic overestimate of the masses, or an underestimate of the luminosities for the bulge giants. However, the 1σ uncertainties on the mean $\Delta(\text{Fe})$ and $\Delta(\text{Ti})$ values at 0.04 dex and 0.06 dex, respectively, indicate that these systematic differences are not statistically significant. Unfortunately, with our results the use of ionization equilibrium to determine stellar masses is only accurate at the factor of 2 uncertainty level and so cannot be used to constrain the average mass of our sample of BW giants.

3.7. Hyperfine Structure

The nuclei of the odd-numbered elements Sc, V, and Co possess nonzero nuclear spin, which results in considerable splitting of the lines into hyperfine components. For all but the most metal-poor of our stars these elements are represented by strong lines only. Because of the desaturating effect of the hyperfine structure, it is absolutely necessary to consider such structure in the abundance analysis of strong lines. This is unlike the case for weak lines, where the hfs merely changes the line profile without altering the equivalent width, leaving the derived abundance unaffected.

We computed the wavelengths of the hfs components using the hyperfine splitting constants A found in the literature. The levels are distinguished by the quantum number F , which takes on values from $I + J$ to $|I - J|$, resulting in $(2J + 1)$ levels in cases where $I \geq J$ or $(2I + 1)$ levels if $J \leq I$. J is the total angular momentum quantum number of the electronic term, and I is the nuclear spin quantum number. The allowed transitions between two terms are $\Delta F = 0, \pm 1$, with $0 \rightarrow 0$ forbidden. The splitting energy is given approximately by equation (5).

$$\Delta E = \frac{A}{2} [F(F + 1) - I(I + 1) - J(J + 1)]. \quad (5)$$

For Sc, V, and Co the ground-state nuclear spin quantum numbers are all $7/2$. The relative strengths of each hyperfine component were found using the table of multiplet line strengths in White & Eliason (1933); to use this table for hyperfine components it was necessary to make the substitutions of F for J , I for S , and J for L . The sum of all components was scaled to the total line gf value as listed in Table 8.

For the Sc I line at 6210.7 Å we used the hfs constants given by Kopfermann & Rasmussen (1934), while the accurate measurements by Arnesen et al. (1982) provided the constants for the Sc II 6604.6 Å line.

The V I splitting constants were not available for every line with an EW measurement; for only those lines at 6812.4,

6274.7, and 6285.2 Å could hfs calculations be performed. The V I hfs data were taken from Childs et al. (1979), Gough et al. (1985), Murakawa (1956), and Kopfermann & Rasmussen (1936). During hfs analysis we found that the V I line at 6812 Å gave significantly higher V abundance values; additionally, the solar line is 4 times stronger than expected. From the Kurucz (1992) line list we determine that this line is most likely contaminated with a Nd II feature, and so is of little use for abundance analysis.

Hyperfine splitting constants for Co I were chiefly taken from Guthöhrlein & Keller (1988, 1990). For the Co I $z^4 D_{5/2}^o$ state we considered the measurements of Oertel (1970), but increased the value of the splitting constants by 15% owing to an apparent offset between the Oertel (1970) hfs constants and

the more reliable measurements of Guthöhrlein & Keller (1988). Unfortunately, no hfs data were available for the Co I $a^2 D_{3/2}$ state involved in the 7417.4 Å transition.

The calculated hfs component wavelengths and gf values are listed in Table 9. These line lists were used in spectrum synthesis calculations to find the best-fit abundance for each line. For lines with EW measurements but no hfs data we used the abundance derived from the EW method only when the line EW was ~ 50 mÅ or less.

The Li line at 6708 Å is, strictly speaking, affected by three forms of hyperfine structure: doublet splitting, isotope splitting, and splitting due to interaction with the nucleus. The wavelengths of these components have been listed by Andersen, Gustafsson, & Lambert (1984). For the purpose of this

TABLE 9
HYPERFINE COMPONENTS

Ion	Wavelength (Å)	gf	Ion	Wavelength (Å)	gf
Co I	6188.932	1.628E-04	Sc I	6210.643	2.522E-03
Co I	6188.946	8.141E-04	Sc I	6210.644	3.363E-03
Co I	6188.978	2.467E-04	Sc I	6210.645	2.589E-03
Co I	6188.990	4.160E-04	Sc I	6210.652	1.681E-03
Co I	6189.003	1.628E-04	Sc I	6210.656	0.000E+00
Co I	6189.017	2.662E-04	Sc I	6210.661	1.615E-03
Co I	6189.026	1.628E-04	Sc I	6210.665	2.522E-03
Co I	6189.038	2.467E-04	Sc I	6210.666	6.672E-03
Co I	6189.048	2.296E-04	Sc I	6210.673	3.363E-03
Co I	6189.055	3.094E-05	Sc I	6210.682	2.589E-03
Co I	6189.064	2.662E-04			
Co I	6189.073	1.449E-04	Sc II	6604.562	2.105E-03
Co I	6189.077	1.628E-06	Sc II	6604.569	3.036E-03
Co I	6189.084	2.296E-04	Sc II	6604.574	7.824E-03
Co I	6189.092	7.978E-05	Sc II	6604.575	2.965E-03
Co I	6189.096	1.449E-04	Sc II	6604.579	3.138E-03
			Sc II	6604.581	1.987E-03
Co I	6678.773	2.653E-04	Sc II	6604.583	6.181E-04
Co I	6678.800	7.960E-04	Sc II	6604.586	2.347E-05
Co I	6678.801	7.576E-04	Sc II	6604.588	1.322E-03
Co I	6678.822	7.960E-04	Sc II	6604.591	2.105E-03
Co I	6678.834	1.669E-03	Sc II	6604.593	3.036E-03
Co I	6678.849	5.691E-04	Sc II	6604.594	1.987E-03
			Sc II	6604.594	2.965E-03
Co I	6814.895	1.192E-03			
Co I	6814.921	1.548E-03	V I	6274.608	1.020E-03
Co I	6814.930	3.071E-03	V I	6274.630	3.061E-03
Co I	6814.943	1.161E-03	V I	6274.642	2.914E-03
Co I	6814.949	7.432E-04	V I	6274.656	6.418E-03
Co I	6814.964	0.000E+00	V I	6274.658	3.061E-03
Co I	6814.976	7.739E-04	V I	6274.679	2.189E-03
Co I	6814.985	1.192E-03			
Co I	6814.993	1.548E-03	V I	6285.108	2.681E-04
Co I	6814.997	1.161E-03	V I	6285.127	7.204E-04
			V I	6285.132	1.977E-03
Co I	7052.842	3.582E-03	V I	6285.144	1.290E-03
Co I	7052.851	2.181E-03	V I	6285.147	2.865E-03
Co I	7052.859	1.135E-03	V I	6285.158	1.935E-03
Co I	7052.889	8.453E-04	V I	6285.159	2.815E-03
Co I	7052.890	1.225E-03	V I	6285.162	8.377E-03
Co I	7052.890	1.203E-03	V I	6285.167	1.935E-03
Co I	7052.891	8.274E-04	V I	6285.172	5.102E-03
Co I	7052.868	4.119E-04	V I	6285.179	2.656E-03
Co I	7052.907	8.274E-04	V I	6285.182	9.634E-04
Co I	7052.914	5.516E-04			
Co I	7052.921	3.080E-04	V I	6812.365	3.674E-03
Co I	7052.928	1.146E-04	V I	6812.370	5.140E-03
			V I	6812.374	4.892E-03
			V I	6812.438	1.077E-02
			V I	6812.444	5.140E-03
			V I	6812.449	1.713E-03

work we adopt a ${}^6\text{Li}/{}^7\text{Li}$ ratio of zero, since Andersen et al. (1984) find an upper limit of 10% in local disk dwarfs. For BW I-194, with a Li EW of 215 mÅ, we find that inclusion of hfs effects reduces the Li abundance from $\epsilon(\text{Li}) = 2.33$ down to $\epsilon(\text{Li}) = 1.72$ dex.

In this paper we find that the hfs calculations reduces derived abundances (relative to the EW method) by as much as 0.7 dex for Co I in μ Leo, and by ~ 0.4 – 0.5 dex for stars at $[\text{Fe}/\text{H}] \sim -0.2$. In Table 10 we compare α Boo abundances derived from the EW method with abundances obtained using hfs calculations.

3.8. Error Analysis

The copious iron lines allow us to estimate uncertainties associated with our abundance results. Figure 8*b* and Table 11 show that the measured dispersion of iron abundances about the mean is typically $\sigma = 0.20$ – 0.25 dex, which we assume to be the uncertainty on an abundance measurement made from a single line. Since we include 60–95 Fe I lines in our analysis, the internal error on the mean Fe I abundance is only 0.02–0.03 dex. As Table 12 shows, the actual $[\text{Fe}/\text{H}]$ uncertainty is larger and is sensitive to the uncertainty in the microturbulent velocity parameter; if $\Delta\xi = 0.3 \text{ km s}^{-1}$, then $\Delta[\text{Fe}/\text{H}] = 0.14$ dex, while for more reasonable microturbulent velocity uncertainties $[\text{Fe}/\text{H}]$ is certain at the 0.10 dex level.

In order to estimate the uncertainties of our abundance analysis, we have constructed Table 12, which demonstrates the effect on derived abundance of errors in the input parameters. From Table 12 we find that the total uncertainty for an abundance ratio, $\Delta[\text{M}/\text{Fe}]$, from a single neutral metal line to $[\text{Fe}/\text{H}]$ (using all Fe I lines) is typically 0.21 dex. This estimate of the ratio uncertainty is based on $\Delta T = \pm 100 \text{ K}$, $\Delta \log g = \pm 0.3$ dex, $\Delta\xi = \pm 0.3 \text{ km s}^{-1}$, $\Delta\text{EW} = \pm 10 \text{ mÅ}$, and $\Delta \log gf = \pm 0.10$ dex. Note that the uncertainty in abundance ratio is less than the quadrature sum of the abundances for two lines, because systematic abundance errors from uncertain atmosphere parameters partially cancel in the ratio due to similar effects for the neutral metal and iron lines.

The abundance ratio uncertainty has random and systematic components. The random errors, arising from EW measurement and adopted $\log gf$ values, are typically $\Delta\epsilon \approx 0.16$ dex, while the systematic errors due to uncertain model atmosphere parameters are $\Delta\epsilon \approx 0.13$ dex. When a species is represented by several lines, the random component of the uncer-

TABLE 10
COMPARISON OF α BOO EW AND HFS ABUNDANCES

Species	Wavelength (Å)	EW (mÅ)	$\epsilon(\text{hfs})$	$\epsilon(\text{ew})$	$\epsilon(\text{ew} - \text{hfs})$
Sc I	6210.671	116	2.95	3.16	0.21
Sc II	6604.600	87	2.87	2.95	0.08
V I	6274.658	130	3.57	3.95	0.38
V I	6285.165	122	3.37	3.63	0.26
V I	6812.356	35	3.75	3.83	0.08
Co I	6188.998	104	4.46	5.26	0.80
Co I	6814.961	121	4.67	5.30	0.63

TABLE 11A
AVERAGE ABUNDANCES FOR BW IV-167

Species	$\epsilon(M)$	$[\text{M}/\text{Fe}]$	σ	N	Notes
Li I	≤ 0.77	≤ -0.83	...	1	ew
Mg I	8.18	0.16	0.15	3	ss
Al I	7.16	0.25	...	1	ss
Si I	7.99	0.00	0.13	6	ss
Ca I	6.75	-0.05	0.13	4	ss
Sc I	4.14	0.60	...	1	hfs
Sc II	3.54	0.00	...	1	hfs
Ti I	5.64	0.21	0.10	9	ss
V I	4.34	-0.10	...	1	hfs
Cr I	6.21	0.10	0.30	2	ew
Fe I	7.96	0.44	0.29	27	ew
Fe II	7.88	0.36	0.03	2	ss
Co I	5.56	0.20	0.01	2	hfs
Ni I	6.60	-0.09	0.10	2	ew
Y II	3.05	0.37	...	1	ew
La II	1.70	0.04	...	1	ew
Eu II	1.05	0.10	...	1	ss

TABLE 11B
AVERAGE ABUNDANCES FOR μ LEO

Species	$\epsilon(M)$	$[\text{M}/\text{Fe}]$	σ	N	Notes
Li I	0.58	-1.00	...	1	ew
Mg I	8.11	0.11	0.07	3	ss
Al I	6.99	0.10	...	1	ss
Si I	8.17	0.20	0.19	6	ss
Ca I	6.52	-0.26	0.07	4	ss
Sc I	3.52	0.00	...	1	hfs
Sc II	3.52	0.00	...	1	hfs
Ti I	5.31	-0.10	0.16	8	ss
V I	3.94	-0.48	0.03	2	hfs
Cr I	6.09	0.00	0.60	2	ew
Fe I	7.94	0.42	0.19	30	ew
Fe II	7.82	0.30	0.14	2	ss
Co I	5.40	0.06	0.02	2	hfs
Ni I	6.82	0.15	0.37	2	ew
Y II	2.71	0.05	...	1	ew
La II	1.59	-0.05	...	1	ew
Eu II	0.63	-0.30	...	1	ss

TABLE 11C
AVERAGE ABUNDANCES FOR BW IV-025

Species	$\epsilon(M)$	$[\text{M}/\text{Fe}]$	σ	N	Notes
Li I	≤ 0.88	≤ -0.65	...	1	ew
Na I	7.25	0.55	...	1	ew
Mg I	8.02	0.07	0.06	3	ss
Al I	7.64	0.80	...	1	ew
Si I	7.67	-0.25	0.20	7	ss
Ca I	7.01	0.28	0.05	4	ss
Sc I	4.07	0.60	...	1	hfs
Sc II	3.67	0.20	...	1	hfs
Ti I	5.77	0.41	0.18	8	ss
V I	4.47	0.10	0.01	2	hfs
Cr I	6.22	0.18	0.17	2	ew
Fe I	7.89	0.37	0.23	46	ew
Fe II	7.45	-0.07	0.24	2	ss
Co I	5.54	0.25	0.22	2	hfs
Ni I	6.59	-0.03	0.22	7	ew
Y II	2.49	-0.12	...	1	ew
La II	1.55	-0.04	...	1	ew
Eu II	0.98	0.10	...	1	ss

TABLE 11D
AVERAGE ABUNDANCES FOR BW IV-072

Species	$\epsilon(M)$	$[M/Fe]$	σ	N	Notes
Li I	≤ 0.30	≤ -0.81	...	1	ew
[O I]	8.58	-0.30	...	1	ss, solar CN mix
Na I	7.24:	0.96	...	1	ew (EW=168 mÅ)
Mg I	8.01	0.48	0.14	3	ss
Al I	7.06	0.64	...	1	ew
Si I	7.56	0.06	0.26	5	ss
Ca I	6.62	0.31	0.19	4	ss
Sc II	3.05	0.00	...	1	hfs
Ti I	5.42	0.48	0.13	9	ss
V I	4.29:	0.34	...	1	hfs (EW=190 mÅ)
Cr I	5.24	-0.38	...	1	ew
Fe I	7.47	-0.05	0.23	35	ew
Fe II	7.27	-0.25	0.27	2	ss
Co I	5.18	0.31	...	1	hfs
Ni I	5.98	-0.22	0.18	4	ew
Y II	2.17	-0.02	...	1	ew
La II	1.13	-0.04	...	1	ew
Eu II	0.86	0.40	...	1	ss

TABLE 11E
AVERAGE ABUNDANCES FOR BW I-322

Species	$\epsilon(M)$	$[M/Fe]$	σ	N	Notes
Li I	≤ 0.10	≤ -0.84	...	1	ew
[O I]	8.71	0.00	...	1	ss, solar CN mix
Na I	6.33	0.22	...	1	ew
Mg I	7.60	0.24	0.16	3	ss
Al I	6.59	0.34	...	1	ew
Si I	7.28	-0.05	0.16	6	ss
Ca I	6.08	-0.06	0.15	4	ss
Sc I	3.13	0.25	...	1	hfs
Sc II	2.98	0.10	...	1	hfs
Ti I	5.00	0.23	0.15	9	ss
Ti II	5.07	0.30	...	1	ew
V I	3.67	-0.11	...	1	hfs
Cr I	5.17	-0.28	...	1	ew
Fe I	7.30	-0.22	0.31	51	ew
Fe II	7.01	-0.51	0.06	2	ss
Co I	4.83	0.13	0.19	2	hfs
Ni I	5.86	-0.17	0.27	7	ew
Y I	1.91	-0.11	...	1	ew
Y II	1.71	-0.31	0.02	2	ew
Zr I	1.66	-0.72	...	1	ew
La II	1.04	0.04	0.05	2	ew
Eu II	0.39	0.10	...	1	ss

TABLE 11F
AVERAGE ABUNDANCES FOR BW I-141

Species	$\epsilon(M)$	$[M/Fe]$	σ	N	Notes
Li I	0.80	-0.13	...	1	ew
Na I	6.20	0.10	...	1	ew
Mg I	7.78	0.43	0.06	3	ss
Al I	7.10	0.86	0.18	2	ew
Si I	7.40	0.08	0.13	6	ss
Ca I	6.16	0.03	0.30	4	ss
Sc I	3.27	0.40	...	1	hfs
Sc II	3.27	0.40	...	1	hfs
Ti I	5.06	0.30	0.13	9	ss
Ti II	4.96	0.20	0.28	2	ss
V I	3.71	-0.06	...	1	hfs
Cr I	5.44	0.00	0.18	2	ew
Fe I	7.29	-0.23	0.24	42	ew
Fe II	7.38	-0.14	0.34	2	ss
Co I	5.04	0.35	...	1	hfs
Ni I	6.04	0.02	0.11	6	ew
Y I	2.01	0.00	...	1	ew
Y II	2.07	0.06	0.08	2	ew
Zr I	1.51	-0.86	...	1	ew
La II	1.00	0.01	0.06	2	ew
Eu II	0.53	0.25	...	1	ss

TABLE 11G
AVERAGE ABUNDANCES FOR BW I-194

Species	$\epsilon(M)$	$[M/Fe]$	σ	N	Notes
Li I	1.72	0.82	...	1	ewhfs
Na I	6.34	0.27	0.37	2	ew
Mg I	7.80	0.48	0.11	3	ss
Al I	6.84	0.63	...	1	ew
Si I	7.43	0.14	0.10	5	ss
Ca I	6.14	0.04	0.21	4	ss
Sc I	3.44	0.60	...	1	hfs
Sc II	3.24	0.40	...	1	hfs
Ti I	5.15	0.42	0.12	9	ss
Ti II	5.25	0.52	0.08	2	ew
V I	3.93	0.19	0.02	2	hfs
Cr I	5.45	0.04	0.17	2	ew
Fe I	7.26	-0.26	0.30	64	ew
Fe II	7.17	-0.35	0.12	2	ss
Co I	4.94	0.28	0.05	2	hfs
Ni I	5.97	-0.02	0.16	8	ew
Y I	2.13	0.15	...	1	ew
Y II	2.22	0.24	...	1	ew
Zr I	1.94	-0.40	...	1	ew
Ba II	1.67:	-0.20	...	1	ew
La II	1.10	0.14	...	1	ew
Eu II	0.50	0.25	...	1	ss

TABLE 11H
AVERAGE ABUNDANCES FOR BW I-012

Species	$\epsilon(M)$	$[M/Fe]$	σ	N	Notes
Li I	0.34	-0.48	...	1	ew
[O I]	8.89	0.30	...	1	ss, solar CN mix
Mg I	7.63	0.39	0.01	2	ss
Al I	7.00	0.87	0.33	2	ew
Si I	7.46	0.25	0.13	7	ss
Ca I	6.04	0.02	0.16	4	ss
Sc I	3.46	0.70	...	1	hfs
Sc II	3.16	0.40	...	1	hfs
Ti I	5.08	0.43	0.12	9	ss
Ti II	5.22	0.57	0.04	2	ew
V I	4.12	0.46	0.04	3	hfs, ew
Cr I	5.33	0.00	0.10	2	ew
Fe I	7.18	-0.34	0.20	62	ew
Fe II	7.21	-0.31	0.05	2	ss
Co I	5.06	0.48	0.05	2	hfs
Ni I	5.89	-0.02	0.16	7	ew
Y I	1.73	-0.17	...	1	ew
Y II	1.76	-0.14	0.03	2	ew
Zr I	1.61	-0.65	...	1	ew
La II	0.99	0.11	0.02	2	ew
Eu II	0.57	0.40	...	1	ss

TABLE 11I
AVERAGE ABUNDANCES FOR BW III-152

Species	$\epsilon(M)$	[M/Fe]	σ	N	Notes
Li I	≤ -0.21	≤ -0.88	...	1	ew
[O I]	8.54	0.10	...	1	ss, solar CN mix
Na I	5.73	-0.11	0.16	2	ew
Mg I	7.51	0.42	0.05	2	ss
Al I	6.79	0.81	0.35	2	ew
Si I	7.70	0.64	0.21	7	ss
Ca I	5.86	-0.01	0.07	4	ss
Sc I	2.91	0.30	...	1	hfs
Sc II	3.31	0.70	...	1	hfs
Ti I	4.69	0.19	0.08	9	ss
Ti II	4.90	0.40	0.40	2	ew
V I	3.35	-0.16	...	1	hfs
Cr I	4.86	-0.32	...	1	ew
Fe I	7.03	-0.49	0.37	65	ew
Fe II	7.09	-0.43	0.09	2	ss
Co I	4.85	0.42	0.02	3	hfs, ew
Ni I	5.76	0.00	0.30	8	ew
Y I	1.27	-0.48	...	1	ew
Y II	2.14	0.39	0.20	2	ew
Zr I	1.51	-0.60	...	1	
Ba II	1.75:	0.11	...	1	ew (EW=194 mÅ)
La II	0.85	0.12	0.01	2	ew
Eu II	0.47	0.45	...	1	ss

tainty is diminished by the square root of the number of lines. With our data the abundance ratio uncertainties for an individual star cannot be improved below the systematic error, of about 0.13 dex, by increasing the number of lines. Based on Table 12 and the number of lines available in Table 11, the uncertainties associated with the [Mg/Fe] abundance ratios are ± 0.17 dex; for [Si/Fe] and [Ca/Fe] we find ± 0.15 dex, while the [Ti/Fe] uncertainty is ± 0.14 dex.

More accurate abundance ratios are possible if individual metal and Fe lines are selected which exhibit similar sensitivities to atmosphere parameters; however, to take advantage of

TABLE 11J
AVERAGE ABUNDANCES FOR BW IV-329

Species	$\epsilon(M)$	[M/Fe]	σ	N	Notes
Li I	0.01	-0.30	...	1	ew
Na I	5.33	-0.15	0.01	2	ew
Mg I	7.27	0.54	0.16	3	ss
Al I	6.22	0.60	0.16	2	ew
Si I	7.06	0.36	0.10	6	ss
Ca I	5.88	0.37	0.20	4	ss
Sc I	2.50	0.25	...	1	hfs
Sc II	2.70	0.45	...	1	hfs
Ti I	4.54	0.40	0.14	9	ew
Ti II	4.39	0.25	0.04	2	ew
V I	3.36	0.15	0.10	3	hfs, ew
Cr I	4.95	0.13	0.26	3	ew
Fe I	6.67	-0.85	0.22	99	ew
Fe II	6.87	-0.65	0.11	2	ss
Co I	4.40	0.33	0.05	3	hfs, ew
Ni I	5.44	0.04	0.23	16	ew
Y I	1.49	0.10	...	1	ew
Y II	1.33	-0.06	0.51	2	ew
Zr I	1.45	-0.30	...	1	ew
Ba II	1.83:	0.55	...	1	ew (EW=190m Å)
La II	0.58	0.21	0.04	2	ew
Eu II	0.14	0.48	...	1	ss

TABLE 11K
AVERAGE ABUNDANCES FOR α BOO

Species	$\epsilon(M)$	[M/Fe]	σ	N	Notes
Li I	≤ -0.82	≤ -1.44	...	1	ew
Na I	5.89	0.10	...	1	ew
Mg I	7.34	0.30	0.04	2	ew
Al I	6.42	0.49	0.13	2	ew
Si I	7.31	0.30	0.16	8	ew
Ca I	5.99	0.17	0.15	5	ew
Sc I	2.95	0.39	...	1	ewhfs
Sc II	2.87	0.31	...	1	ewhfs
Ti I	4.77	0.32	0.13	8	ew
Ti II	4.66	0.21	0.23	2	ew
V I	3.56	0.10	0.19	3	ewhfs
Cr I	5.10	-0.03	0.23	4	ew
Fe I	6.98	-0.54	0.18	107	ew
Fe II	6.95	-0.57	0.22	3	ew
Co I	4.57	0.19	0.15	2	ewhfs
Ni I	5.84	0.13	0.23	15	ew
Y I	1.16	-0.54	...	1	ew
Y II	1.28	-0.42	0.30	2	ew
Zr I	0.99	-1.07	...	1	ew
Ba II	1.75	0.16	...	1	ew (EW=161 mÅ)
La II	0.53	-0.15	0.01	2	ew
Eu II	0.29	0.32	...	1	ew

line abundance ratios with very small systematic errors, it would be necessary to make accurate EW measurements to reduce the random errors, which would require very high S/N spectra. The systematic component of the abundance ratio errors could also be reduced by averaging the [M/Fe] ratios over several stars.

4. DISCUSSION

4.1. Standard Stars

A comparison of our [Fe/H] results for α Boo and μ Leo with those of Branch, Bonnell, & Tomkin (1978) and Sneden & Gratton (1990) appears in Table 13; note that we adopt

TABLE 11L
AVERAGE ABUNDANCES FOR BW IV-003

Species	$\epsilon(M)$	[M/Fe]	σ	N	Notes
Li I	0.46	0.24	...	1	ew
[O I]	8.09	0.10	...	1	ss, solar CN mix
Mg I	6.93	0.29	0.16	2	ss
Al I	5.70	0.17	...	1	ew
Si I	6.88	0.27	0.16	5	ew
Ca I	5.85	0.43	0.34	7	ew
Sc II	2.46	0.30	...	1	hfs
Ti I	4.40	0.35	0.10	9	ew
Ti II	4.57	0.52	0.43	2	ew
V I	2.98	-0.08	0.43	2	ew
Cr I	4.87	0.14	0.18	3	ew
Fe I	6.58	-0.94	0.22	87	ew
Fe II	6.65	-0.87	0.01	2	ss
Co I	4.23	0.25	0.13	4	ew
Ni I	5.36	0.05	0.18	18	ew
Y I	1.80:	0.50	...	1	ew
Y II	1.35	0.05	0.43	2	ew
Zr I	1.98:	0.32	...	1	ew
Ba II	1.52	0.33	...	1	ew
La II	0.51	0.23	0.01	2	ew
Eu II	-0.18	0.25	...	1	ss

TABLE 11M
AVERAGE ABUNDANCES FOR BW IV-203

Species	$\epsilon(M)$	$[M/Fe]$	σ	N	Notes
Li I	-0.46	-0.54	...	1	ew
[O I]	7.85	0.00	...	1	ss, solar CN mix
Na I	5.08	-0.17	...	1	ew
Mg I	6.84	0.34	0.23	2	ss
Al I	5.84	0.45	0.28	2	ew
Si I	6.96	0.49	0.27	7	ew
Ca I	5.45	0.17	0.10	5	ew
Sc I	2.07	0.05	...	1	hfs
Sc II	2.22	0.20	...	1	hfs
Ti I	4.19	0.28	0.11	3	ew
Ti II	4.51	0.60	0.33	2	ew
V I	2.88	-0.04	0.03	3	ew
Cr I	4.54	-0.05	0.19	4	ew
Fe I	6.44	-1.08	0.23	80	ew
Fe II	6.60	-0.92	0.22	2	ss
Co I	3.91	0.07	0.09	4	hfs, ew
Ni I	5.19	0.02	0.23	14	ew
Y I	0.82	-0.34	...	1	ew
Y II	1.27	0.11	0.36	2	ew
Zr I	1.01	-0.51	...	1	ew
La II	0.36	0.22	0.09	2	ew
Eu II	-0.47	0.10	...	1	ss

NOTE TO TABLES 11A-11M.—Columns from left to right indicate, respectively, species identification, average abundances $\epsilon(M)$ on a logarithmic scale of $H = 12.00$, logarithmic ratios $[M/Fe]$ relative to the solar values, standard deviation of abundances, number of lines, and remarks concerning synthesis method.

KEY TO NOTES.—ew—equivalent width matching calculations; ss—spectrum synthesis profile matching including CN lines and Kurucz 1992 atomic lines; hfs—spectrum synthesis profile matching with hyperfine structure components used in the line list; ewhfs—equivalent width matching calculations using hyperfine structure components only; solar CN mix—indicates that the solar C/Fe and N/Fe ratios were adopted in the molecular equilibrium calculations.

$\epsilon(Fe)$ for the Sun as 7.52 on the $H = 12.00$ scale. For the Fe I results at least our abundance scale is in excellent agreement with these previous studies; this is particularly pleasing for our μ Leo results.

Abundances for other elements are not always in such good agreement. In μ Leo the important α -element ratios $[Mg/Fe]$, $[Si/Fe]$, $[Ca/Fe]$, and $[Ti/Fe]$ found here are 0.11, 0.20, -0.26, and -0.10 dex, respectively, whereas SG90 find -0.15, 0.01, -0.23, and -0.14, respectively. The Ca and Ti abundances are in good accord, but there is conflict at the ± 0.2 dex level for Si and Mg. The disparity in $[Mg/Fe]$ ratios is not serious, as the SG90 result is based on a single line. Abundances of other elements in μ Leo agree with SG90 at the 0.2 dex level, with the exception of V, for which SG90 obtained a larger abundance by 0.36 dex.

Our α Boo abundances are in excellent agreement with the recent work of Peterson et al. (1993), although this is not surprising because we used the Griffin (1968) Arcturus atlas as the source of equivalent widths. Peterson et al. (1993) find $[Fe/H] = -0.5 \pm 0.1$, with α -element ratios $[Mg/Fe]$, $[Si/Fe]$, $[Ca/Fe]$, and $[Ti/Fe]$ at 0.4, 0.4, 0.3, and 0.3, respectively. In this work we find $[Fe/H] = -0.54$ and α -elements at 0.30, 0.30, 0.17, and 0.32, respectively. For other species there is excellent agreement between this work and Peterson et al. (1993), with differences up to about 0.1 dex.

The good agreement found with these recent analyses suggests that our EW measurement procedure works well even in stars as metal-rich as μ Leo, and that our abundance analysis code is correct. The differences between our Mg and Si abundances in μ Leo and the results of SG90 might indicate a problem with the spectrum profile matching procedures.

We find it satisfying that our μ Leo abundance for the important element Ti is in accord with the result of SG90; this supports our Ti results for the BW giants.

4.2. Iron Abundances

The $[Fe/H]$ values found in this paper are somewhat less than found for the same stars by Rich (1988). Rich (1988)

TABLE 12
DEPENDENCE OF ABUNDANCES ON INPUT PARAMETERS

Ion	λ (Å)	χ (eV)	EW (mÅ)	$\Delta\epsilon(M)$				$[M/Fe]^a$
				ΔT +100 K	$\Delta\xi$ +0.3 km s ⁻¹	$\Delta\log g$ +0.3 cm s ⁻²	ΔEW +15 mÅ	
Mg I ^b	6319.2	5.11	...	+0.03	-0.06	+0.03	+0.11 ^c	0.16
Al I	6698.7	3.14	90	+0.09	-0.10	-0.01	+0.22	0.19
Si I	7415.9	5.61	88	-0.07	-0.08	+0.08	+0.22	0.21
Ca I	6166.4	2.52	133	+0.13	-0.18	-0.01	+0.21	0.21
Ca I	6508.8	2.52	67	+0.12	-0.09	-0.01	+0.24	0.21
Ti I	6861.5	2.27	112	+0.17	-0.16	+0.01	+0.22	0.22
Eu II	6645.1	1.38	46	0.00	-0.07	+0.11	+0.27	0.24 ^d
Ba II	6496.9	0.60	158	+0.05	-0.29	+0.07	+0.21	0.28 ^d
Fe II	6432.7	2.89	42	-0.11	-0.10	+0.15	+0.31	...
$[Fe/H]^e$	...	...	...	+0.05	-0.12	+0.05	...	...

^a Quadrature sum of abundance ratio errors, with $\Delta EW = 10$ mÅ and $\Delta\log gf = 0.10$ dex.

^b Spectrum synthesis of Mg I feature with a depth of 24%.

^c Increase in abundance if depth is increased by 2%.

^d Relative to Fe II rather than Fe I abundance.

^e Average iron abundance $[Fe/H]$ from all lines with $EW \leq 150$ mÅ.

TABLE 13
COMPARISON OF [Fe/H] WITH OTHER WORK^a

Star	BBT	SG ^b	MR
α Boo	-0.49	-0.45	-0.54
μ Leo	+0.48	+0.45	+0.42

^a BBT: Branch, Bonnell, & Tomkin 1978; SG: Sneden & Gratton 1990; MR: this work.

^b We have corrected the SG abundance scale to a solar $\epsilon(\text{Fe}) = 7.52$.

gave [Fe/H] results for 88 BW stars, based on a variety of abundance indicators; his preferred method (solution 1) relied upon Fe I and Mg I features, while his solution 2 relied upon Fe I features only. On the average our [Fe/H] values are 0.32 dex smaller than Rich's solution 1 values and 0.25 dex smaller than his solution 2 results. In Figure 15 we plot [Fe/H] found here with the R88 solution 1 results for all 11 stars. It is clear from Figure 15 that the R88 [Fe/H] values are in reasonably good agreement at low metallicity only. A linear regression fit provides the following transformation between the R88 and iron abundances found in this paper:

$$[\text{Fe}/\text{H}](\text{R88}) = 0.381 + 1.190[\text{Fe}/\text{H}](\text{MR}) \quad (6)$$

This regression is represented by the dashed line in Figure 15; the rms dispersion about the relation is 0.18 dex, in accord with R88's noise estimate of 0.16 dex.

Since the preferred R88 [Fe/H] index uses a combination of Mg and Fe features, it is in fact not a true measure of [Fe/H] but rather of $[(\text{Mg} + \text{Fe})/\text{H}]$. The assumption by R88 that $[\text{Fe}/\text{H}] = [(\text{Mg} + \text{Fe})/\text{H}]$ can only be valid if the Mg/Fe ratio follows the relation for the disk/halo standard stars used by R88. However, our high-resolution analysis shows that the BW giants are enhanced in Mg/Fe relative to the solar value for all but the most metal-rich stars. The good agreement between R88 solution 1 [Fe/H] values seen in Figure 15 for stars near $[\text{Fe}/\text{H}] \sim -1$ suggests that the magnesium abundances may be responsible for the divergence seen in Figure 15.

We can make a very crude comparison of the R88 $[(\text{Mg} + \text{Fe})/\text{H}]$ index with our abundances by combining our Mg and

Fe abundances in the appropriate proportions, using equation (7):

$$[(\text{Mg} + \text{Fe})/\text{H}] = \log_{10} [(1 - f) \times 10^{[\text{Fe}/\text{H}]} + f \times 10^{[\text{Mg}/\text{H}]}], \quad (7)$$

where

$$f = \frac{\text{EW}(\text{Mg } b)}{\text{EW}(\text{Fe I } \lambda 5270) + \text{EW}(\text{Fe I } \lambda 5335) + \text{EW}(\text{Mg } b)}. \quad (8)$$

The variable f simply measures the fractional contribution of Mg to the R88 Mg + Fe index, with EWs as given in R88. The variable f was typically about 0.35 and ranged from 0.21 to 0.46. The resultant $[(\text{Mg} + \text{Fe})/\text{H}]$ values from this work are compared with R88 solution 1 in Figure 16. Figure 16 shows excellent agreement between our Mg + Fe index and the R88 index, with a rms deviation of 0.25 dex. It is reasonable to conclude that R88's use of Mg + Fe lines as a general abundance indicator may have led to the high derived abundances for bulge stars. Furthermore, in this work we have seen the importance of molecular blanketing on the derived Fe abundance for our high-resolution, high-S/N analysis. It must be even more difficult to determine accurate abundances for metal-rich stars using the low-resolution data of R88.

In Figure 17 we present a frequency histogram plot of the R88 [Fe/H] results, solution 1 values, corrected by the transformation of equation (6). The corrected R88 solution 1 [Fe/H] values in Figure 17 have a mean of -0.25 and a standard deviation of 0.40 dex, significantly below the R88 values of 0.08 and 0.48 dex, respectively. Our new mean [Fe/H] for the R88 BW giants is 0.11 dex above the mean [Fe/H] for solar neighborhood G dwarfs given by Pagel & Patchett (1975) (mean [Fe/H] = -0.36, $\sigma = 0.26$ dex), but is 0.08 dex below the mean [Fe/H] of 671 solar neighborhood GK giants found by McW90 (mean = -0.17, $\sigma = 0.16$ dex). It should be noted, however, that McW90 used a magnitude-limited sample, including subgiants, giants, and supergiants of spectral types G and K, although only a handful of subgiants and supergiants were observed. Because it was a magnitude-limited sample, the McW90 survey was biased in favor of luminous, more massive than average, giant stars, which would favor slightly higher abundances. Another difficulty for the McW90 survey was the

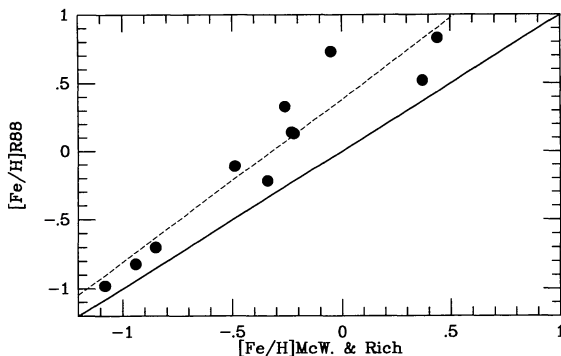


FIG. 15.—Plot of Fe abundances found here with those of Rich (1988), solution 1. The dashed line represents a least-squares regression fit to the points.

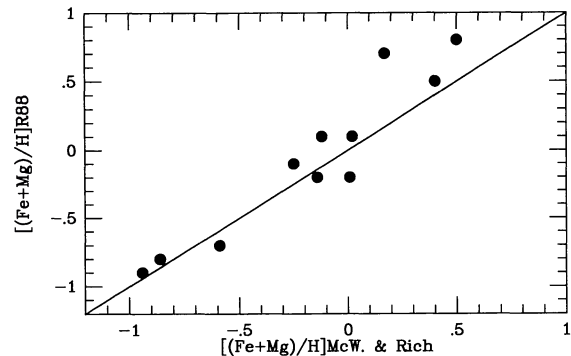


FIG. 16.—Plot of synthetic $[(\text{Fe} + \text{Mg})/\text{H}]$ abundances found here against $[(\text{Fe} + \text{Mg})/\text{H}]$ from Rich (1988), solution 1. The points are well correlated below solar metallicity.

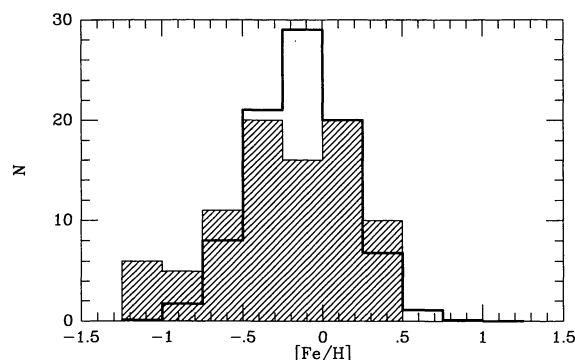


FIG. 17.—The hatched area represents a frequency histogram of Rich (1988) solution 1 $[\text{Fe}/\text{H}]$, corrected for the regression relation found here. The solid histogram outline represents the solar neighborhood GK giant distribution found by McWilliam (1990), smoothed by a Gaussian of $\sigma = 0.25$ dex. Note that the bulge giants have a similar mean $[\text{Fe}/\text{H}]$ but a slightly broader distribution and more very metal-poor stars.

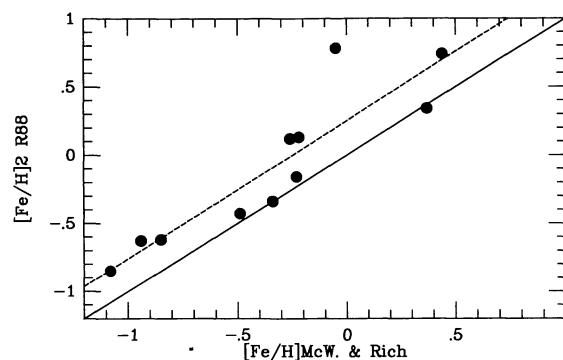


FIG. 18.—Plot of Fe abundances found here with those of Rich (1988) solution 2. The dashed line represents a least-squares regression fit to the points.

absence of metal-rich model atmospheres, which resulted in low $[\text{Fe}/\text{H}]$ estimates for the metal-rich end of the sample. These two deficiencies have opposing effects on the final $[\text{Fe}/\text{H}]$ frequency distribution function, and so must cancel to some degree. Given these reservations, it is still compelling to compare the brightest GK giants in the solar neighborhood with the brightest K giants in the bulge.

In order to compare the McW90 survey with the corrected R88 bulge abundances, we first binned the McW90 results into steps of 0.05 dex in $[\text{Fe}/\text{H}]$. The resultant distribution function was convolved with a Gaussian of $\sigma = 0.25$ dex, rebinned into steps of 0.25 dex, and then scaled to the number of stars in the R88 sample. Note that we convolve the McW90 $[\text{Fe}/\text{H}]$ distribution with an error function broader than the estimated R88 uncertainties; we do this in order to draw conservative conclusions about the bulge $[\text{Fe}/\text{H}]$ function. The two distributions are shown in Figure 17. It is clear that the bulge and solar neighborhood $[\text{Fe}/\text{H}]$ distributions are very similar, although the BW sample contains more stars of low metallicity than the solar neighborhood, and the BW distribution appears to be broader than in the solar neighborhood. The excess of metal-poor stars in the bulge sample may be a manifestation of the so-called G dwarf problem (Rich 1990a) seen in the disk; another possibility is that these stars belong to an extension of the halo population into BW.

We repeat the comparison with R88 $[\text{Fe}/\text{H}]$ results, but using the R88 solution 2 $[\text{Fe}/\text{H}]$ metallicities. Since the solution 2 metallicities used two Fe I lines, rather than two Fe I lines and one Mg I line, the comparison with our $[\text{Fe}/\text{H}]$ values is appropriate. However, since only two lines were used by R88 for this indicator, the dispersion is expected to be worse than the solution 1 measurements. In Figure 18 we present a plot of $[\text{Fe}/\text{H}]$ derived here with the solution 2 $[\text{Fe}/\text{H}]$ of R88. We see that the R88 metallicities are simply shifted by 0.25 dex relative to this work. The formal solution to the regression relation is given in equation (9):

$$[\text{Fe}/\text{H}](\text{MR}) = 0.99[\text{Fe}/\text{H}](\text{R88 solution 2}) - 0.25. \quad (9)$$

The rms dispersion of the regression relation in equation (9) is

0.25 dex. The 0.25 dex shift seen in Figure 18 is somewhat surprising, in light of the good agreement between R88 solution 1 and our $[(\text{Mg} + \text{Fe})/\text{H}]$ index in Figure 16.

In Figure 19 we present a frequency histogram of R88 solution 2 $[\text{Fe}/\text{H}]$ corrected for the regression relation. The corrected R88 $[\text{Fe}/\text{H}]$ data have a mean of -0.21 with $\sigma = 0.49$ dex. In Figure 19 we also indicate the solar neighborhood $[\text{Fe}/\text{H}]$ function, convolved with a Gaussian error function of $\sigma = 0.35$ dex. We see from Figure 19 that the R88 BW $[\text{Fe}/\text{H}]$ distribution is broader than that for the local giants but that the means of the distributions are very similar.

The conservative conclusions to be drawn from the $[\text{Fe}/\text{H}]$ frequency histograms of Figures 17 and 19 are that the corrected R88 metallicity function has a mean value in good agreement with the mean value for the solar neighborhood giants, and that the dispersion in $[\text{Fe}/\text{H}]$ is larger for the R88 BW giants, due to an excess of metal-poor giants. If the observed rms $[\text{Fe}/\text{H}]$ dispersions (from Figs. 15 and 18) between this work and R88 solutions 1 and 2 are used to smooth the McW90 $[\text{Fe}/\text{H}]$ distribution, then an additional conclusion can be drawn: that the R88 BW abundances show a higher frequency of SMR stars in the bulge than in the solar neighborhood.

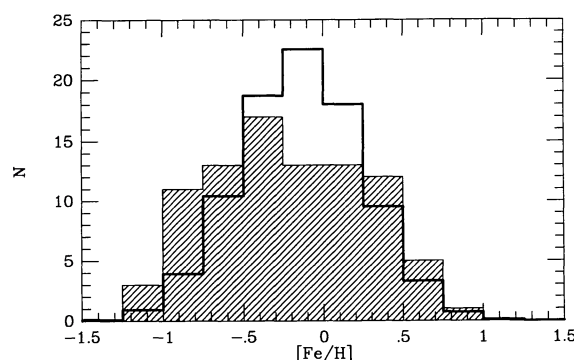


FIG. 19.—The hatched area represents a frequency histogram of Rich (1988) solution 2 $[\text{Fe}/\text{H}]$, corrected for the regression relation found here. The solid histogram outline represents the solar neighborhood GK giant distribution found by McWilliam (1990), smoothed by a Gaussian of $\sigma = 0.35$ dex. We again see that the bulge giants have a similar mean $[\text{Fe}/\text{H}]$ but a slightly broader distribution and more very metal-poor stars.

Furthermore, our analysis of one of the most SMR BW stars in the R88 sample (BW IV-167) indicates that it is only 0.02 dex more metal-rich than the SMR disk giant μ Leo.

It should be recalled that for the most metal-rich stars we found that $(J - K)$ and $(V - K)$ temperatures differ by up to 350 K, and that the hotter, $(J - K)$, temperatures were adopted in the analysis. If the adopted T_{eff} values are reduced downward toward the $(V - K)$ temperatures, the iron abundances derived will decrease, roughly according to the amounts indicated in the error budget table (Table 12). Therefore, the difficulty with IR color temperatures cannot be used to explain the difference between $[\text{Fe}/\text{H}]$ found here and the work of Rich (1988).

Our low Fe abundances relative to R88 solution 2 could be explained by a zero-point error in our adopted temperature scale, caused by an underestimation of the reddening to BW. From Table 12 it is seen that for a 0.25 dex shift in Fe abundance a revision of our temperatures upward by 500 K is necessary. Such an increase in temperature could be accomplished by assuming extra reddening of (using Table 5) $\Delta E(B - V) = +0.15$ for $(J - K)$ temperatures, and $\Delta E(B - V) = +0.22$ for $(V - K)$ temperatures. However, the uncertainty in $E(B - V)$ for BW is relatively small; the 1σ dispersion in BW $E(B - V)$ from the results of Walker & Mack (1986) and Walker & Terndrup (1991) are both 0.05 mag. These results appear to exclude reddening as a source for a 500 K shift in our temperature scale. To increase our iron abundances by 0.25 dex, we could appeal to a zero-point error in the observed broad-band colors used to derive T_{eff} , but the size of the zero-point shifts required are unreasonably large: for $(J - K)$ temperatures a 0.16 mag shift is needed to account for the 500 K, while for $(V - K)$ a zero-point error of 0.50 mag is required. Additionally, zero-point errors in our temperature scale in excess of 150 K are ruled out by the infrared flux temperatures found for FW87 K giants, as discussed in § 3.2.

Our low $[\text{Fe}/\text{H}]$ scale is in discord with the results of Geisler & Friel (1992) and Tyson (1991). Both of these studies used Washington photometry to obtain $[\text{Fe}/\text{H}]$ for GK giants in Baade's window. The Washington photometric system has recently been recalibrated by Geisler, Clariá, & Minniti (1991). For solar metallicity and below, the calibration uses the photometric indices of bright solar neighborhood giants and published $[\text{Fe}/\text{H}]$ abundances from high-resolution spectra. Calibration of the system into the SMR regime utilizes metallicities inferred from DDO photometry. Because the Tyson (1991) work relied upon the R88 zero point, it does not qualify as an independent measurement of the bulge metallicity. An implicit assumption inherent in the use of the Washington system to determine $[\text{Fe}/\text{H}]$ in bulge giants is that the bulge stars possess a mixture of elements with approximately the same recipe seen in the solar neighborhood. It is common for observers to use a measured index to "read off" the $[\text{Fe}/\text{H}]$ value of corresponding local giants. However, if the mixture of elements is unlike that of the local stars, the index most probably will not give the correct Fe abundance, although it could reflect better the overall metal abundance. Such difficulties can be magnified by the blanketing effects of diatomic molecules in SMR stars, because the strength of the molecular absorption increases with the product of the atomic abundances; CN band strength, for example, is very sensitive to the nitrogen abundance. Our abundance analysis of 11 bulge giants indicates

that the chemical composition of the bulge stars differs in several respects from the solar abundance distribution. To avoid future errors in photometric $[\text{Fe}/\text{H}]$ values, the Washington system should be calibrated for the Galactic bulge using Galactic bulge standards, which have been thoroughly examined with high-resolution abundance analysis. In future it will probably be wise for photometrists to state what assumptions are made about relative abundances of the elements when quoting metallicity results.

As discussed earlier, the apparent difference between $(V - K)$ and $(J - K)$ temperatures may be indicative of V -band blanketing. If such blanketing has occurred, then the metallicities derived from the Washington system could also be affected.

At first sight our low $[\text{Fe}/\text{H}]$ abundance scale also appears to be inconsistent with the IR study of BW M giants of FW87. In their paper FW87 show that the spectral types of M giants are much later than those of field giants at the same $(J - K)_0$. Also, the CO indices of BW M giants are larger than those of field giants. These observations seem to suggest that the metallicity of BW M giants is considerably higher than that of the field giants. We note that the spectral type of M giants is sensitive to the TiO strength, which is governed by T_{eff} and the Ti and O abundances.

The CO indices in M giants are sensitive to the C and O abundances, but because the CO bands are saturated, the abundance sensitivity is comparable to the sensitivity for other parameters (McWilliam & Lambert 1984). Although the calibration of the CO index by McWilliam & Lambert (1984) is specific to the Kitt Peak National Observatory filter set, we can use it to estimate the required abundance and/or atmospheric parameter differences between bulge observations of FW87 and the disk M giants. To account for the 0.03 mag shift in the FW87 CO indices, systematic shifts in a number of parameters could be invoked: an increase in the mean BW M giant carbon abundance of 0.2 dex (assuming solar C/O ratio), a 0.4 dex decrease in $\log g$, a 0.4 km s^{-1} increase in the mean microturbulence, or a change in mean $^{12}\text{C}/^{13}\text{C}$ ratio from 25 to 7. Obviously, numerous combinations of shifts in these parameters could be chosen to make up for the 0.03 mag shift in mean CO index. In disk stars there is an anticorrelation between gravity and microturbulent velocity, such that the luminous, low-gravity stars possess higher microturbulent velocity parameters. In particular, if the BW M giant masses are lower, on average, than the observed solar neighborhood giants, we can expect the lower mean gravity and concomitant high microturbulent velocities to lead to a CO index increase which could easily be interpreted as an enhanced carbon abundance. We are not trying to repudiate the suggestion of FW87 that the systematically high CO index seen in BW M giants is an abundance effect, but wish to make it clear that an interpretation of this observation is somewhat clouded.

Our abundance results are qualitatively consistent with the difference between TiO spectral type in BW and disk M giants observed by FW87, since our Ti/Fe ratios are in excess of the solar value for all program stars. For stars near the mode of the R88 abundance distribution, our abundance results indicate that $[\text{Ti}/\text{Fe}] \sim 0.4$ dex. The relatively low abundance of Ti (the solar Fe/Ti ratio being ~ 340) makes it an insignificant contributor to the overall metal abundance, even at $[\text{Ti}/\text{Fe}] \sim 0.4$ dex.

FW87 (and Frogel et al. 1984) pointed out that the infrared H-R diagram [K_0 versus $(J - K)_0$ plot] of the BW M giants is in apparent contradiction to the supposed SMR nature of these stars, because the median line of the BW giant branch runs close to the giant branch of 47 Tuc. In fact, FW87 show that the red edge of the BW giant branch is similar to, but slightly redder than, the M67 giant branch. It was suggested by FW87 that an unidentified blanketing source caused the $(J - K)$ colors to be bluer than anticipated for SMR bulge stars.

Infrared photometry of the BW giant-branch colors was extended an additional 4 mag, well into the K giant realm, by Davidge (1991). This new photometry confirms FW87's finding that the BW colors approximate those of 47 Tuc. We conjecture that any unknown blanketing effect (most likely due to molecular absorption) sufficient to cause bluer $(J - K)$ colors in the M giants would have abated at the higher temperature of the K giants, due to diminished molecular formation, and the K giant branch would be observed redder, moving toward that expected for a SMR giant branch. The absence of a red, SMR, BW K giant branch supports the notion that BW giants are not extremely metal-rich, although age must also enter into the interpretation of the giant branch. We note that if some unknown blanketing has affected the K giant $(J - K)$ colors, it would be necessary to decrease the temperatures of our BW sample (particularly for our most metal-rich stars), resulting in *even lower* iron abundances (by perhaps 0.25 dex) than presently obtained. Finally, the correlation found for $(J - K)$ and IR flux temperatures seen in § 3.2 indicates that the putative molecular blanketing is nonexistent.

We believe that the apparent contradiction between BW M giant spectral type- $(J - K)$ color relation (supporting a SMR bulge), and the BW K_0 - $(J - K)$ plot (supporting a "metal-normal" bulge) is partly resolved by our observations of relatively normal iron abundances, at $[\text{Fe}/\text{H}] = -0.25$, and a supersolar Ti/Fe ratio in the bulge K giant stars. Given a mean BW metallicity of $[\text{Fe}/\text{H}] = -0.25$, the position of the BW giant branch in the infrared H-R diagram of FW87 requires greater masses for the BW giants than the globular cluster giants. FW87 claim that the BW giant branch is coincident with the 47 Tuc locus; if this is true, then the BW giants must be more massive than the 47 Tuc giants, because our mean BW $[\text{Fe}/\text{H}]$ is 0.60 dex larger than recent estimates of $[\text{Fe}/\text{H}]$ for 47 Tuc (e.g., $[\text{Fe}/\text{H}] = -0.85$ found by Brown & Wallerstein 1991).

Given the observation (FW87) that the BW giants lie on the 47 Tuc giant-branch locus, the mass difference between the BW giants and the 47 Tuc giants can be estimated if the $\Delta T/\Delta \text{Mass}$ and $\Delta T/\Delta \text{Fe}$ gradients for these red giant stars are known. We first attempted a mass estimate for the BW giants using the observed properties of 47 Tuc and M67 RGB loci relative to the BW giant branch, as shown in the IR H-R diagram of FW87. Employing the assumption that metallicity scales with $[\text{Fe}/\text{H}]$, and using the M67 $[\text{Fe}/\text{H}]$ of 0.04, a turn-off mass of $1.24 M_\odot$ (Garcia Lopez, Rebolo, & Beckman 1988), and the theoretical tracks of Maeder & Meynet (1988), we found a mean giant mass of $1.35 M_\odot$. This mass estimate is well in excess of even the highest recent values indicated by age estimates for the bulge of 8–9 Gyr (Holtzman et al. 1993). A second calculation of the mean BW giant mass used the theoretical evolutionary tracks of Vandenberg (1992) to estimate the $\Delta T/\Delta \text{Mass}$ and $\Delta T/\Delta \text{Fe}$ gradients; this also gave an approximate mean mass of $1.35 M_\odot$.

If the $[\text{Fe}/\text{H}]$ difference between 47 Tuc and BW is in fact only 0.50 dex, as would be allowed by measurement uncertainty, and we choose the smallest gradients from the two methods used above, then our estimate of the mean BW giant mass can be brought down to $1.20 M_\odot$. A further reduction in mass is allowed if it is assumed that the BW $[\text{O}/\text{Fe}]$ ratio is solar, but $[\text{O}/\text{Fe}] = 0.40$ dex in 47 Tuc. Using Vandenberg's (1992) approximate rule, that a 0.25 dex increase in $[\text{O}/\text{Fe}]$ leads to an age decrease of ~ 1 Gyr, we can assign a slightly generous correction of $0.08 M_\odot$, resulting in a minimum mean mass near $1.1 M_\odot$ for BW giants.

The apparently high luminosities of BW giants could also be explained if the adopted BW distance is too large. The adopted FW87 BW distance of 7 kpc is shorter than many literature estimates, which tend to put the bulge distance closer to 8 kpc. In the unlikely event that the Galactic center is 6 kpc rather than 7 kpc distant, then the luminosities of giants in the FW87 IR H-R diagram should be reduced by ~ 0.3 mag. Such a shift would reduce our mean mass estimate for BW giants from ~ 1.4 to about $1.1 M_\odot$.

We conclude that, based on the FW87 H-R diagram and our mean BW $[\text{Fe}/\text{H}]$, the mean BW giant mass is probably not lower than $1.1 M_\odot$, although the best estimate, $1.35 M_\odot$, appears to be unacceptably large. We note that if Baade's window is actually more distant than 7.0 kpc, our estimates of the mean giant mass will increase, at approximately $0.16 M_\odot$ for every 0.5 kpc.

Our low $[\text{Fe}/\text{H}]$ scale is further supported by the initial results of an abundance analysis of the putative SMR Galactic bulge M giants (Smith 1993). Smith finds that the handful of so called SMR M giants actually range in $[\text{Fe}/\text{H}]$ from -0.2 to 0.2 , and are on average about 0.2 dex lower in $[\text{Fe}/\text{H}]$ than claimed by Cook (1987).

Due to the controversial nature of our low $[\text{Fe}/\text{H}]$ results for BW, we performed one final check of our abundance scale. The check involved abundance analysis of a red giant, star 563 (Menzies 1974), in the globular cluster NGC 5927. Six frames, with a total of 2 hr integration time, of star 563 were acquired during the 1991 CTIO bulge run; the S/N of the resultant spectrum was about 30. Model atmosphere parameters ($T_{\text{eff}} = 4400$ K, $\log g = 1.7$) were taken from Frogel, Persson, & Cohen (1983). Equivalent widths for the cleanest 26 Fe I lines in the spectrum were used to derive $[\text{Fe}/\text{H}] = -0.38 \pm 0.07$. Full details of the analysis will be published elsewhere (McWilliam & Rich 1994). This abundance is in excellent agreement with the scale of Zinn & West (1984), who obtained $[\text{Fe}/\text{H}] = -0.30 \pm 0.09$ for the cluster. Friel & Geisler (1991) find $[\text{Fe}/\text{H}] = -0.25 \pm 0.25$, based on Washington photometry, while Armandroff & Zinn (1988) found $[\text{Fe}/\text{H}] = -0.38$ from integrated light spectroscopy of the Ca II infrared triplet. The agreement of our result with the low-resolution and broad-band studies casts considerable doubt on François's (1991) determination of $[\text{Fe}/\text{H}] = -1.08$ from an echelle spectrum of star 532 in NGC 5927. For this cluster R88 derived $[\text{Fe}/\text{H}] = -0.02 \pm 0.15$, and used this result to compare with the BW giants. It is clear that the R88 scale is higher than the confluence of $[\text{Fe}/\text{H}]$ results for NGC 5927 by about 0.3 dex, and this confirms our finding that the BW giant $[\text{Fe}/\text{H}]$ values are ~ 0.25 dex lower than found by R88. The consistency of our NGC 5927 and BW $[\text{Fe}/\text{H}]$ results indicates that our BW metallicities have not been significantly affected by selection

biases. Our result for NGC 5927 is useful also, as it provides a link between our abundance scale for Galactic bulge giants with metallicity determinations of the globular cluster system.

It has been the practice of a few observers of the Galactic bulge (e.g., R88; GF91; Tyson 1991) to convert observed $[\text{Fe}/\text{H}]$ values into Z , the mass fraction of metals. In making this conversion, these authors have once again assumed that all elements vary in lockstep with Fe. However, in this work we demonstrate that the elemental abundance distribution in bulge K giants is not like that of the solar neighborhood. Frequency histograms of Z have often been used to distinguish between simple theoretical models of the nucleosynthetic history of the bulge.

In calculating Z we must know the abundances of all elements, or at least all elements which contribute significantly to the mass fraction of metals. Unfortunately, in our sample of K giants it is rather difficult to derive abundances for the most abundant elements, namely, C, N, O, and Ne. Without reliable estimates of abundances for these elements it is not possible to compute useful Z -values. For example, if the solar oxygen abundance were to be raised by 0.40 dex (as might be expected from a Type II supernova (SN) nucleosynthetic origin) then $\Delta \log Z$ would be increased by 0.24 dex. If C, O, and Ne were each increased by 0.40 dex, then $\Delta \log Z$ would be increased by 0.31 dex. We cannot use other so-called α -elements in the bulge giants to guide our selection of O and Ne abundances, since two α -elements (Mg and Ti) are overabundant and two (Si and Ca) appear normal.

This sensitivity of Z to the most abundant elements is further complicated by CO molecular formation, which produces an interdependence between the measured C and O abundances. In addition, possible CNO nucleosynthetic processing of envelope material, which is seen in many luminous cool giants, could mask the primordial O abundance. Future studies of bulge giants should attempt to evaluate C, N, and O abundances with C_2 features near 5000 Å, $[\text{O I}]$ at 6300 Å, and CN red system lines near 7000 Å. If possible, the target stars should be of relatively low luminosity, to avoid the possibility of nonprimordial O abundances.

It is clear that the current data are insufficient to compute Z for each of the program stars. Furthermore, it is not even reasonable to scale the solar O and Ne abundance to elements with measured abundances in our stars.

4.3. The Galactic $[\text{Fe}/\text{H}]$ Gradient

The results of our abundance analysis impact on the Galactic $[\text{Fe}/\text{H}]$ gradient, although a comparison of bulge with disk K giants may be inappropriate, and a steady trend from one component to the other is not necessarily expected. However, if the Galactic $[\text{Fe}/\text{H}]$ gradient seen in the solar neighborhood is extended to the bulge, then average disk stars at the edge of the bulge will have $[\text{Fe}/\text{H}] \sim 0.5$ dex larger than our sample of bulge stars. This would imply that at the edge of the bulge there is a discontinuous step down in $[\text{Fe}/\text{H}]$ from super-metal-rich disk stars to a solar neighborhood metallicity bulge population. To investigate this unusual possibility, we briefly consider the evidence for a Galactic $[\text{Fe}/\text{H}]$ gradient.

The iron abundance gradient in the Galaxy has been studied by numerous authors: a photometric *UBV* analysis of

~ 100 K giants toward the Galactic center by Helfer & Jennens (1975) found no difference in average $[\text{Fe}/\text{H}]$ for K giants along the line of sight toward the Galactic center. Approximate distances of these K giants from the Sun ranged from 0.1 to 6.5 kpc.

Janes (1979), using DDO photometry of distant K giants and *UBV* photometry of open clusters, found an overall $[\text{Fe}/\text{H}]$ gradient of -0.05 dex kpc^{-1} and found that the trend steepens beyond a Galactocentric distance of about 11 kpc. The CN strengths of stars toward the Galactic center were fairly constant, which led Janes to suggest that the gradient in the inner disk is “perhaps zero.”

Luck (1982) used G and K supergiants in the range $7.7 \text{ kpc} \leq R \leq 10.2 \text{ kpc}$ (assuming $R_\odot = 8.5 \text{ kpc}$) to find a gradient $d[\text{Fe}/\text{H}]/dR = -0.13 (\pm 0.03)$ dex kpc^{-1} .

Shaver et al. (1983) used radio and optical data of H II regions to measure the gradients of O, N, S, Ne, and Ar with Galactocentric radius, reaching 50% of the distance to the Galactic center. For O and N they found gradients of -0.07 and -0.09 dex kpc^{-1} , respectively; but for S and Ne they found very shallow gradients, at -0.01 dex kpc^{-1} . Our oxygen abundances would suggest a flat gradient within the solar circle, in disagreement with the Shaver et al. O gradient; however, we note that our oxygen abundances were derived under the assumption of solar C/Fe and N/Fe ratios. We also note the possibility that our sample of stars may have undergone oxygen depletion through the CNO cycle processing, rather like the globular cluster giants of Sneden et al. (1991). Unfortunately, our abundance analysis did not include N, S, Ne, or Ar, so further comparison with the Shaver et al. work is not possible. A discrepancy between H II region and K giant abundance ratios is not of great concern; we point out that H II regions trace the Galactic gas, whereas a sample K giant stars is composed of material spanning many gigayears of Galactic history.

Neese & Yoss (1988) studied 364 G and K stars, using DDO and *BV* photometry in the Galactic center and anticenter directions; Galactocentric radius ranged from ~ 4 to 14 kpc. They found an overall $[\text{Fe}/\text{H}]$ gradient of $-0.017 (\pm 0.007)$ dex kpc^{-1} , but their Figure 6 showed that a zero $[\text{Fe}/\text{H}]$ gradient is completely consistent with their data for stars within the solar circle.

In a critical study of literature iron abundances Janes, Tilley, & Lyngå (1988) studied various properties of the open cluster system. Their Figure 8 showed a plot of $[\text{Fe}/\text{H}]$ against Galactocentric distance, ranging from 6 to 14 kpc, which indicated an overall $[\text{Fe}/\text{H}]$ gradient of about -0.07 dex kpc^{-1} . However, the figure clearly shows that for open clusters within a Galactocentric radius of 10 kpc a zero (or near-zero) $[\text{Fe}/\text{H}]$ gradient provides a good description of the data.

A study of giants in old Galactic clusters by Friel & Janes (1993) claims that the old and young clusters give a similar $[\text{Fe}/\text{H}]$ gradient, at -0.09 ± 0.02 dex kpc^{-1} . As with previous cluster studies, most of the points lie beyond the solar circle. Although the Friel & Janes (1993) data within the solar circle appear to follow the regression of $[\text{Fe}/\text{H}]$ with distance in the anticenter direction, an inspection of their Figure 2 indicates that the data within 10 kpc of the center are also consistent with a zero $[\text{Fe}/\text{H}]$ gradient. We contend that none of the studies of the Galactic $[\text{Fe}/\text{H}]$ gradient have clearly demon-

strated a nonzero gradient within the solar circle, and that extant data can be described with a zero $[\text{Fe}/\text{H}]$ gradient. Resolution of this question will have to await studies of disk clusters closer to the Galactic center.

The small gradient in $[\text{Fe}/\text{H}]$ out to the solar circle, followed by a marked decrease in $[\text{Fe}/\text{H}]$ of the outermost regions, is in good qualitative agreement with model 9 of Larson (1976). The steeper gradient in the outer regions occurs in large part because star formation proceeds slowly, so that much of the material is gaseous and not yet processed by stars.

4.4. *Alpha Elements*

The abundances of the so-called α -elements in our stars is of particular utility as a probe into the formation timescale of the Galactic bulge. In the Galactic disk and halo the α -elements (O, Mg, Si, Ca, and Ti) follow a trend of increasing abundance, relative to Fe, as $[\text{Fe}/\text{H}]$ decreases from the solar value to about -1.0 (e.g., Lambert 1987; Wheeler, Sneden, & Truran 1989). Current popular thought maintains that this is due to the delay between nucleosynthesis by high-mass, Type II SNe and low-mass Type I SNe.

Nucleosynthesis calculations (e.g., Arnett 1971, 1978; Woosley 1986; Hillebrandt 1978) indicate that Type II SNe produce large amounts of oxygen, other α -elements, r -process elements, and some iron; Type I SN events, on the other hand, produce large quantities of iron-peak elements and little else. These calculations indicate that the efficiency of oxygen and other α -element production increases with increasing initial mass of the progenitor.

Since Type II SNe are more massive than Type I progenitors, the Type II SNe are responsible for the initial nucleosynthesis, and produce abundance ratios rich in α - and r -process elements relative to iron. When the first Type I SNe begin to detonate, the $[\alpha/\text{Fe}]$ ratio starts to decline toward solar values. The timescale for the transition point of the $[\alpha/\text{Fe}]$ ratio to begin decreasing depends upon the mass of Type I progenitors and is approximately 10^8 – 10^9 yr (e.g., Wheeler et al. 1989).

Therefore, if the Galactic bulge formed quickly, on a timescale of 1 Gyr or less, one would anticipate high $[\alpha/\text{Fe}]$ ratios to be found in most bulge stars. This qualitative expectation is borne out by a more detailed quantitative model of the Galactic bulge formation (Matteucci & Brocato 1990). We point out that for the elements Si, S, Ca, and Ti it cannot be stated that the SN nucleosynthesis calculations indicate that their abundances should vary in lockstep with oxygen. It is the observations of the α -element abundances in the disk and halo stars which suggest that this may be true. The simple picture of chemical evolution of the α -elements relies on the assumption that the initial mass function (IMF) allows formation of significant numbers of stars in the Type II and Type I SN mass range. The 1 Gyr timescale limit could be extended indefinitely, by simply allowing the IMF to restrict formation of stars in the Type I progenitor mass range for an arbitrary period of time.

The $[\alpha/\text{Fe}]$ ratios from Table 11 are plotted against $[\text{Fe}/\text{H}]$ in Figures 20a–20d. One can immediately see that the emerging picture is not uniform across the α -elements: Mg and Ti remain at ~ 0.4 dex for all but the two most metal-rich stars, while Si and Ca behave very much like the stars in the Galactic

disk/halo. Furthermore, the oxygen abundances derived from the 6300 Å [O I] line suggest that the mean $[\text{O}/\text{Fe}]$ is solar at almost all $[\text{Fe}/\text{H}]$, with a possibility that $[\text{O}/\text{Fe}]$ declines for metal-rich stars. The oxygen abundance result for BW IV-003 is in severe disagreement with an earlier preliminary result of $[\text{O}/\text{Fe}] \sim 1$ (McWilliam & Rich 1990). The 1990 result was based on our 1990 spectrum for BW IV-003 only, for which $\text{S/N} \sim 26$. We suspect that the high $[\text{O}/\text{Fe}]$ value resulted, in part, from the poor S/N. Also, the [O I] EW and continuum points were measured by hand (not using an automated routine such as GETJOB), and we suspect that the continuum was placed too high.

Under the ad hoc assumptions that Mg and Ti behave alike and that Si behaves as Ca, we present Figure 20c, in which averages of $[\text{Mg}/\text{Fe}] + [\text{Ti}/\text{Fe}]$ and $[\text{Si}/\text{Fe}] + [\text{Ca}/\text{Fe}]$ are plotted against $[\text{Fe}/\text{H}]$. It is somewhat frustrating to have Si and Ca suggest that the bulge nucleosynthesis was dominated by Type I SNe over a time interval exceeding 1 Gyr, while Mg and Ti suggest that formation was less than 1 Gyr and nucleosynthesis dominated by Type II SNe. If all of our abundance ratios are correct, then we must conclude that reality is more complex than the simple picture of α -element synthesis outlined above.

Note that the differences between Mg, Ti and Si, Ca abundance ratios are considerably larger than the abundance ratio uncertainty of about ± 0.15 dex. Furthermore, Table 12 can be used to show that it is not possible to modify the atmosphere parameters in such a way that all four α -elements give the same $[\alpha/\text{Fe}]$ ratio. The small scatter between $[\alpha/\text{Fe}]$ ratios within our sample of stars further suggests that the differences seen in Figures 20a–20d for Mg, Ti and Si, Ca reflect genuine abundance discrepancies.

It would be fascinating to understand the α -element abundances with detailed SN nucleosynthesis calculations. Potentially, it would be possible to identify the chemistry with mean SN mass. Current predictions of isotopic abundance ratios from O to Ni fit the solar abundance distribution remarkably well, with an uncertainty of ~ 0.3 dex over a range of seven orders of magnitude in abundance ratio. Unfortunately, the 0.3 dex uncertainty is the approximate scale of our abundance variations, so the theoretical predictions are not yet accurate enough for a detailed comparison with observations. We suggest that our α -element abundance ratios are not impossible to attain. For example, Mg produced in hydrostatic burning lies in a region between 1.8 and 6.0 $M_\odot$ from the center (where $\text{Mg}/\text{Si} \sim 10$) in the 25 $M_\odot$ model of Woosley & Weaver (1982), and is not affected by shock-induced nucleosynthesis. In the same model, hydrostatically synthesized Si lies between 1.5 and 1.8 $M_\odot$ from the center, while the postshock Si lies between 1.7 and 1.9 $M_\odot$ and has been significantly affected by explosive nucleosynthesis. It is clear, as Woosley & Weaver (1986) point out, that the synthesis of Si through Ca is somewhat sensitive to the energy of the shock. Arnett (1978) states: “Because they probe a deep layer of the star, near or at the mass cut, the abundances of Si to Fe provide a critical test of models of supernova explosions and gravitational collapse.” These considerations lead us to believe that our unusual Mg/Si and Mg/Ca abundance ratios are not ruled out by SN models but, if correct, might provide useful constraints on SN models.

We remind the reader that our Mg and Fe abundances indi-

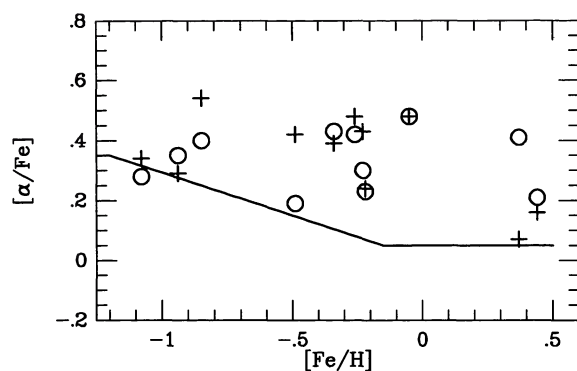


FIG. 20a

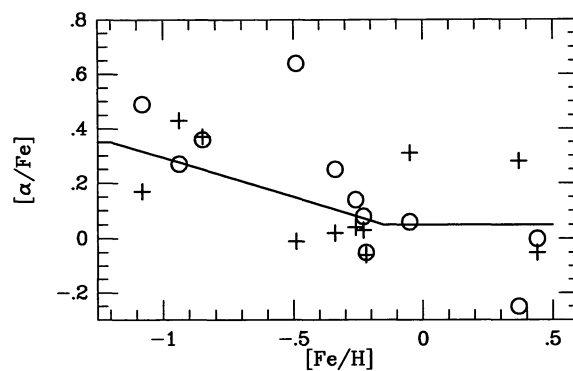


FIG. 20b

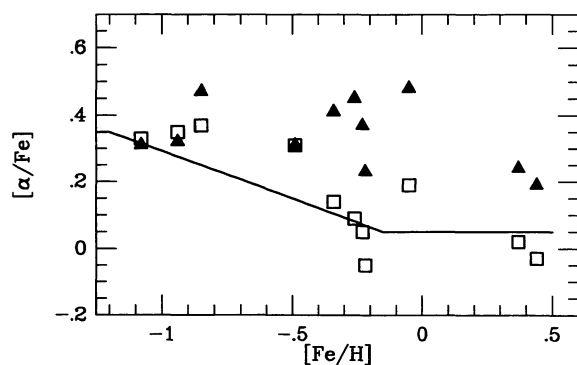


FIG. 20c

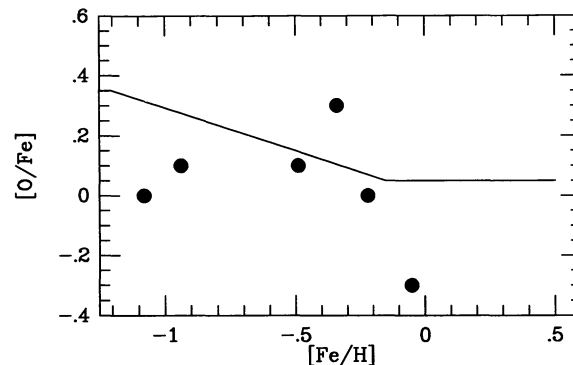


FIG. 20d

FIG. 20.—(a) Plot of $[\alpha/\text{Fe}]$ against $[\text{Fe}/\text{H}]$ for our sample of 11 bulge giants. Open circles represent $[\text{Ti}/\text{Fe}]$, and plus signs represent $[\text{Mg}/\text{Fe}]$. The solid line is the disk/halo relation found by Lambert (1987). (b) Plot of $[\alpha/\text{Fe}]$ against $[\text{Fe}/\text{H}]$ for our sample of 11 bulge giants. Open circles represent $[\text{Si}/\text{Fe}]$, and plus signs represent $[\text{Ca}/\text{Fe}]$. The solid line is the disk/halo relation found by Lambert (1987). (c) Plot of $[\alpha/\text{Fe}]$ against $[\text{Fe}/\text{H}]$ for our sample of 11 bulge giants. Filled triangles represent an average of $[(\text{Mg} + \text{Ti})/\text{Fe}]$, while open squares indicate the average $[(\text{Si} + \text{Ca})/\text{Fe}]$. The solid line is the disk/halo relation found by Lambert (1987). (d) Plot of $[\text{O}/\text{Fe}]$ against $[\text{Fe}/\text{H}]$ for six bulge giants with measurable $[\text{O I}]$ features. The solid line is the disk/halo relation found by Lambert (1987). Note that the sample can be accurately described with $[\text{O}/\text{Fe}] = 0.0$.

cate that for most of the bulge giants the Mg/Fe ratio is well in excess of the solar value. High Mg abundances are not without precedent; studies of luminous elliptical galaxies (Worthey et al. 1992) claim that the *average* $[\text{Mg}/\text{Fe}] \approx 0.2\text{--}0.3$ in these systems. In terms that mirror our findings for the bulge giants, they suggest that “the well known metal enhancement in giant ellipticals is due almost entirely to overabundance of the light elements and that the Fe-peak itself is only slightly enhanced.” It is noteworthy that Worthey et al. and R88 measure virtually the same features: Mg , $\text{Fe I } \lambda 5270$, and $\text{Fe I } \lambda 5335$. Before concluding that we can generalize our findings about the bulge stars to the giant elliptical galaxies, it is important to recall that, while bulges are in general supported by rotation, an anisotropic stellar velocity dispersion supports the shape of ellipticals (Kormendy & Illingworth 1982). It is reasonable to speculate that ellipticals have even shorter formation and enrichment timescales than bulges.

The Ti I lines are unusually strong in our sample of BW stars, a fact which might be used to argue that the abundances are unduly sensitive to upper atmosphere effects and should not be used for abundance analysis. In Figures 21a and 21b we compare EWs of Ti I lines in BW bulge stars near $[\text{Fe}/\text{H}] =$

-0.25 with disk K giants of the same $[\text{Fe}/\text{H}]$ and $\log g$, taken from McW90. The figures demonstrate that the BW Ti I EWs are larger than those seen in disk K giants of the same $[\text{Fe}/\text{H}]$, $\log g$, and T_{eff} , providing qualitative proof of a higher Ti abundance in the BW stars. Only an unacceptably large systematic error in T_{eff} (of about 450 K) would be capable of bringing the BW Ti I EWs into line with those seen in disk K giants (which would reduce the $[\text{Fe}/\text{H}]$ values even more). We note that our abundance analysis for μ Leo gave a Ti result in good agreement with the much higher resolution study of SG90; this confirmation suggests that the high $[\text{Ti}/\text{Fe}]$ ratios found for the BW stars are correct.

High $[\text{Ti}/\text{Fe}]$ ratios can resolve our normal $[\text{Fe}/\text{H}]$ result with the high metallicities suggested by FW87, which they based partly on the bulge M giant spectral type–($J - K$) relation. The assumption implicit in the high metallicities suggested by FW87 was that the M giant spectral type depends on metallicity; although this can be true, more specifically the spectral type is a function of TiO band strength, which in normal circumstances is largely controlled by the Ti abundance. Therefore, we take the FW87 observation of late spectral type–($J - K$) relation as supporting evidence for our high Ti abun-

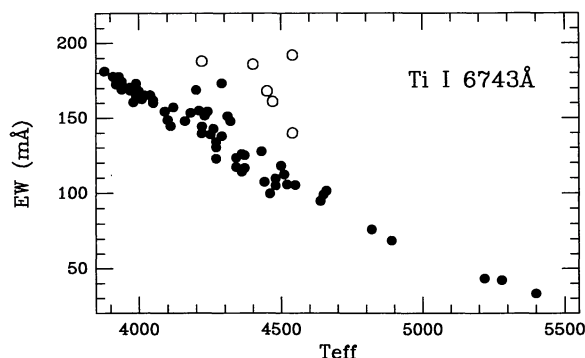


FIG. 21a

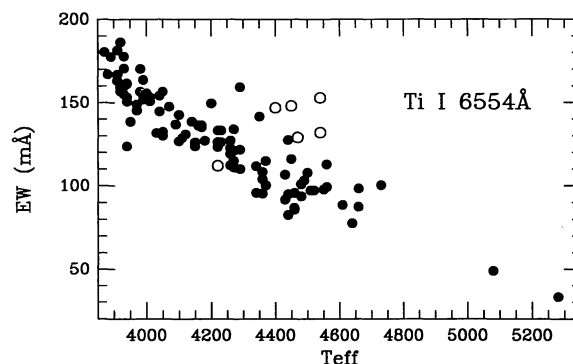


FIG. 21b

FIG. 21.—(a) Plot of equivalent width of the Ti I line at 6743 Å against effective temperature for stars in the solar neighborhood sample (filled circles) of McWilliam (1990), and the six BW bulge giants (open circles) with $[\text{Fe}/\text{H}]$ near -0.25 . Selection from the McWilliam (1990) sample has been restricted to stars with $-0.13 \geq [\text{Fe}/\text{H}] \geq -0.27$ and $2.6 \geq \log g \geq 1.4$. To shift the bulge points onto the locus of the local giants would require a temperature reduction of ~ 450 K. (b) Same as (a), but for the Ti I line at 6554 Å. The low bulge point is for BW I-322. Notice that this line also requires a temperature shift of ~ 400 K to move the bulge stars onto the locus of local giants.

dances. We note that even large increases in the Ti abundance will not significantly affect overall stellar metallicities, since Ti is relatively rare compared to other metals.

The solar $[\text{O}/\text{Fe}]$ ratios seen for nearly all stars in Figure 20d are at odds with expectations based on the disk/halo $[\text{O}/\text{Fe}]$ trend with iron abundance for “normal” stars. The relatively low oxygen abundances could be understood in the context of the luminous globular cluster stars at the tip of the giant branch as seen by Sneden et al. (1991), McWilliam, Geisler, & Rich (1992), and Drake, Smith, & Suntzeff (1992). Some of these giants possess low O/Fe ratios compared to the remainder of giants in each cluster. Also, it is seen that these giants possess varying, but huge, overabundances of Al (e.g., Drake et al. 1992; McWilliam et al. 1992), characteristic of nuclear processing by proton burning of the stellar envelope material. In our sample of BW giants, which are relatively luminous, most possess large Al abundances, and one star is super-Li-rich, in accord with nuclear processing of the envelope. Until more detailed and extensive studies are performed, including a set of low-luminosity bulge giants, we cannot rule out the possibility that the oxygen abundances have been severely depleted by CNO processing. We must therefore regard oxygen as being of questionable use at the present time.

We conclude that our pattern of α -element abundances is unlike that seen in the disk or the halo of the Galaxy. Therefore, the history of the stars responsible for nucleosynthesis in the bulge must have differed from the disk or halo. If we assume that the Si and Ca abundances have been affected by different SN nucleosynthesis in the bulge, and rely upon the Mg and Ti abundances as α -element indicators, we would conclude that almost all of the bulge nucleosynthesis was performed by Type II SNe, and that the formation time for the bulge was quite short, under 1 Gyr.

If we were to draw a conclusion taking our oxygen abundances at face value, we would suggest that Type I SNe contributed significantly to the nucleosynthesis of bulge material, and that the formation timescale involved was greater than 1 Gyr. This question must be resolved by further observations, including stars of lower luminosity which should be unaffected by depletion of primordial oxygen.

4.5. Heavy Elements

The s -process elements in our analysis are represented by Y I, Zr I, Y II, Ba II, and La II lines. In order to investigate the trend of s -process abundance with metallicity, it is best to ratio ionized lines with the Fe II abundance and neutral lines with the Fe I abundance. In this way errors due to uncertainties in the stellar atmosphere parameters will be canceled to first order. Of the s -process lines measured, the Zr I line appears to be systematically low relative to the other lines, and the Ba II line is frequently too strong for a reliable abundance determination; we therefore omit using Zr I abundances, and only use the Ba II abundance for two stars.

From the lines of ionized s -process elements we compute a mean overabundance relative to the Fe II abundance for each star. This $[s/\text{Fe}]$ was usually derived from the Y II and La II lines, but in two cases the Ba II abundance could be included. In Figure 22 we plot the quantity $[s/\text{Fe}]$ against $[\text{Fe}/\text{H}]$; we also include the $[\text{Y I}/\text{Fe I}]$ ratios, which are identified with crosses on the plot. It is clear from Figure 22 that the ratio s/Fe is largely independent of $[\text{Fe}/\text{H}]$, at approximately $[s/\text{Fe}] \sim$

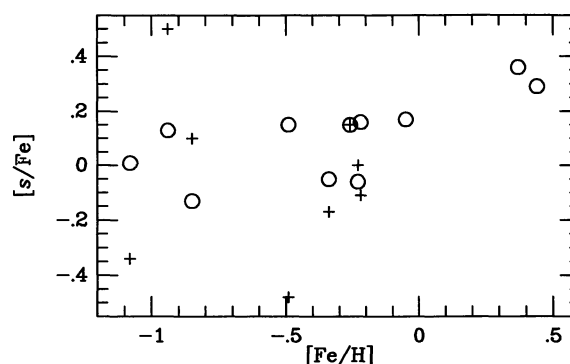


FIG. 22.—Plot of mean s -process abundance, $[s/\text{Fe}]$, against $[\text{Fe}/\text{H}]$ for our sample of bulge giants. Open circles represent the mean overabundance of $[(\text{Y II} + \text{La II} + (\text{Ba II}))/\text{Fe II}]$, and plus signs represent $[\text{Y I}/\text{Fe I}]$. The $[s/\text{Fe}]$ trend with $[\text{Fe}/\text{H}]$ is well described by $[s/\text{Fe}] = 0.0$, identical to the disk/halo relation of Lambert (1987). The most metal-rich stars may be slightly enhanced in s -process abundances.

0.0. There also seems to be a slight increase in $[s/\text{Fe}]$ for the two most metal-rich stars (BW IV-4167 and BW IV-4025). Because of the high metallicity of these two stars, it is plausible that this small upturn in $[s/\text{Fe}]$ is simply a consequence of unknown contaminating features affecting the equivalent widths of the s -process lines.

The constant value of $[s/\text{Fe}]$ at $[s/\text{Fe}] \sim 0.0$ is reminiscent of the disk and halo systems of the Galaxy at $[\text{Fe}/\text{H}] \geq -1.5$; below $[\text{Fe}/\text{H}]$ of about -1.5 , the s/Fe abundance ratio declines rapidly with decreasing iron abundance. Our $[s/\text{Fe}]$ result does not impact on the question of Type II or Type I nucleosynthesis sites, but we do find some significance in the fact that $[s/\text{Fe}]$ is not markedly lower than the solar ratio, as is seen in the oldest, most metal-poor, halo stars.

Abundances for the only available r -process element, Eu, are plotted in Figure 23; the Eu II abundances are ratioed with iron abundances derived using the available Fe II lines. It is apparent from this figure that $[\text{Eu II}/\text{Fe II}]$ is higher than the ratio for disk giants in two of the metal-rich stars (BW IV-025 and BW IV-072). These high Eu values would indicate greater nucleosynthesis from Type II SNe than is seen in the disk stars, for, as with the α -elements, Eu is synthesized almost exclusively in Type II SNe. However, because the Fe II abundances are based on only two or three lines (depending on blanketing), we give this result with some caution. The reader may notice that the mean of the Eu distribution for giants in Figure 23 does not intersect the solar value. The source of the discrepancy could be non-LTE overpopulation of the 6645 Å lower level, a saturation effect in the McW90 analysis which did not include hfs synthesis abundances, or possibly an error in the laboratory gf value for the 6645 Å line (the solar gf value is 0.12 dex larger than the laboratory value). Since the Eu line EWs are relatively small, we do not suspect that the problem lies with the hyperfine splitting. A zero-point shift due to gf -value error, or due to non-LTE effects, should be the same in both bulge and solar neighborhood giants.

We strongly suggest that follow-up studies explore Eu (and other r -process elements like Dy) in solar metallicity bulge giants; higher resolution, greater wavelength coverage, and higher S/N would increase the number of lines for which r -

process abundances could be extracted. Such data would also increase the number of Fe II lines available.

4.6. Light Elements

The high Li abundance found for BW I-194, at $\epsilon(\text{Li}) = 1.7$, confirms our initial suspicion that this star is a super-lithium-rich giant.

A study of 644 GK giants by Brown et al. (1989) indicates that the frequency of super-lithium-rich giants is at the 1% level in the solar neighborhood. It is remarkable that our study of only 11 bulge giants has uncovered such an unusual star, and this could be interpreted as an indication of fundamental differences between bulge and disk giants. Unfortunately, the 1σ uncertainty associated with a single event is 100%, so until at least one more Li-rich bulge giant is found, we cannot state that the frequency of Li-rich stars in the bulge differs from the solar neighborhood frequency.

The origin of high Li abundances in GK giants has been addressed to some extent by Brown et al. (1989). The preferred scenario for Li production is through a hot-bottomed convective envelope (Cameron 1955; Cameron & Fowler 1971), where ${}^7\text{Be}$ is produced via the ${}^3\text{He}(\alpha, \gamma){}^7\text{Be}$ reaction in convective cells which reach deep into the interior of the star, exposing the material to temperatures greater than about 10^7 K. This is followed by a transport of the ${}^7\text{Be}$ to regions with temperatures less than 3×10^6 K and capture of an electron via ${}^7\text{Be}(e, \nu){}^7\text{Li}$, to produce lithium. Any electron capture in regions where the temperature exceeds 3×10^6 K yields lithium which is quickly destroyed by other nuclear reactions. Investigations into this form of Li production have been undertaken by Wallerstein & Sneden (1982) and by Scalo, Despain, & Ulrich (1975, hereafter SDU75). Li production via this mechanism is restricted to evolved stars on the asymptotic giant branch (AGB), with much higher luminosities ($M_{\text{bol}} \sim -5.4$; SDU75) and lower surface temperatures (characteristic of cool M giants) than the GK giants studied by Brown et al. or in this study; SDU75 also provide a lower total mass limit of $\sim 1.5 M_{\odot}$, which is slightly higher than our assumed bulge turnoff mass. This difficulty with the relatively early evolutionary state of the GK Li-rich stars is reminiscent of the problem which existed for Ba II stars (a group of GK giants with high s -process abundances) prior to the discovery by McClure, Fletcher, & Nemec (1980) and McClure (1983) that all Ba II stars possess white dwarf companions. The Ba II stars are now thought to have received mass from a highly evolved AGB companion star, which had previously synthesized the s -process elements during He-shell flashing events. If the Li-rich GK giants are produced like the Ba II stars, then they are mass transfer objects and should each possess a white dwarf remnant of the original primary.

SDU75 show that the hot-bottomed convective envelope would also be responsible for CNO processing. This CNO processing might be invoked to explain the low oxygen abundances found in the bulge giants.

We find $[\text{Al}/\text{Fe}]$ ratios in excess of the solar ratio for all our bulge giants, but particularly in those stars with $[\text{Fe}/\text{H}] \geq -0.5$, for which the mean $[\text{Al}/\text{Fe}]$ ratio is 0.7 dex. BW IV-167 and BW I-322 are slightly anomalous in this group at $[\text{Al}/\text{Fe}] = 0.25$ and 0.34 dex, respectively. The $[\text{Al}/\text{Mg}]$ ratio is not

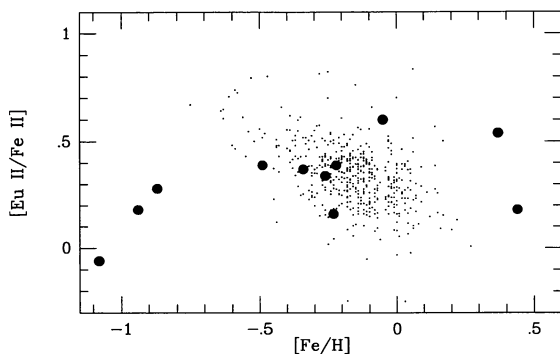


FIG. 23.—Plot of $[\text{Eu II}/\text{Fe II}]$ for the sample of BW bulge giants (large filled circles), compared with the distribution found by McWilliam (1990) (small points) for local GK giants. Two metal-rich stars, BW IV-072 and BW IV-025, appear to be significantly enhanced. The McWilliam (1990) results did not include hfs calculations, which may explain the slight difference between the two samples near $[\text{Fe}/\text{H}]$ of -0.3 .

as well behaved as $[\text{Al}/\text{Fe}]$, but this could be due to increased observational scatter resulting from the use of Mg instead of Fe abundances. If our $[\text{Al}/\text{Mg}]$ ratios are compared to the trend of Lambert (1987), and the theoretical Type II SN yields from Arnett (1971) and Woosley & Weaver (1982), we would anticipate that $[(\text{C} + \text{N} + \text{O})/\text{H}] \sim 0.5$ for many of our program stars. These predicted high C + N + O values, based on an extrapolation of the $[\text{Al}/\text{Mg}]$ trend for SN nucleosynthesis, may be incorrect if we believe that the approximate oxygen abundances found here are of primordial origin.

We note that high $[\text{Al}/\text{Fe}]$ ratios have been seen by numerous authors (e.g., Drake et al. 1992; McWilliam et al. 1992) in some of the most luminous globular cluster giants. There appears to be a correlation of $[\text{Al}/\text{Fe}]$ with CN strength (Norris et al. 1981) and an anticorrelation with oxygen abundance (Drake et al. 1992). Our relatively normal $[\text{Al}/\text{Fe}]$ ratio for BW IV-167 does not fit the CN-Al trend seen by Norris et al. (1981), and the remaining bulge giant Al abundances seem to be independent of CN strength. The Al peculiarities seen in globular clusters are often accompanied by high Na abundances, and there appears to be a real dispersion in both elements for stars in a single cluster (e.g., Kraft et al. 1992). Such abundance peculiarities could be due to proton burning of envelope material, but it is not yet clear that our K giants are in an evolutionary phase capable of producing the observed Al enhancements though single-star evolution. If the high $[\text{Al}/\text{Fe}]$ ratios are not of primordial origin, then it would seem most likely that proton capture occurred in an evolved primary, which subsequently lost mass to the current primary. The observed high $[\text{Al}/\text{Fe}]$ ratios for almost all of our sample and the association of high $[\text{Al}/\text{Fe}]$ with binary mass transfer would suggest that almost all objects in our sample are binaries with white dwarf companions.

If we maintain that the high Al abundances are of primordial origin, then the high $[\text{Al}/\text{Mg}]$ ratios require a modification of SN nucleosynthesis theory, because such high ratios are not predicted in the Arnett (1971) or Woosley & Weaver (1982) supernova nucleosynthesis calculations. We are unable to make any definitive conclusions about the origin of the high Al abundances, but we do note that Peterson et al. (1993) and this work both find a large $[\text{Al}/\text{Fe}]$ ratio for α Boo, which, with its high $[\alpha/\text{Fe}]$ ratios, is similar to stars in the halo. Assuming that α Boo is not a mass transfer object, this observation might argue in favor of a primordial origin for the high Al abundances. The reader should be aware that proton capture nucleosynthesis will not necessarily be seen in iron-peak elements, owing to the large Coulomb repulsion force.

In our sample of BW stars the $[\text{Na}/\text{Mg}]$ ratios appear to follow the relation seen in the disk, as outlined by Lambert (1987). However, the most metal-rich stars exhibit a relatively high ratio, at $[\text{Na}/\text{Mg}] \sim 0.5$.

4.7. Iron Peak

Among our iron-peak elements, V, Cr, and Ni scale closely with $[\text{Fe}/\text{H}]$, showing no obvious trend with metallicity. However, scandium and cobalt appear to be overabundant; the mean $[\text{Co}/\text{Fe}]$ ratio is 0.28 dex. We divide Sc into $[\text{Sc I}/\text{Fe I}]$ and $[\text{Sc II}/\text{Fe II}]$, for which we find the mean values to be 0.42 and 0.33 dex, respectively. The standard deviation of the abun-

dance ratio distributions is 0.20 dex for Sc I, Sc II, V I, and Cr I, and 0.10 dex for Ni I and Co I abundances.

Enhancements of Co and Sc could be indicative of very light neutron exposure of the atmospheric material. As shown by Smith & Lambert (1987), the addition of, on average, one or two neutrons to Fe nuclei can greatly enhance the Co abundance. Because Fe is typically ~ 400 times more abundant than Co, the cobalt abundance can change significantly with negligible effect on the iron abundance. The same principle applies to the addition of neutrons to Ca nuclei. Smith & Lambert predict that such mild *s*-processing is accompanied by modest enhancements of Ba.

5. CONCLUSIONS

We have acquired high-resolution, high-S/N spectra of 12 Galactic bulge K giants previously studied by Rich (1988). From abundance analysis of these spectra and recalibration of the R88 metallicity scale, we find the mean BW $[\text{Fe}/\text{H}]$ to be -0.25 dex, slightly lower than, but not appreciably different from, the solar neighborhood K giants. We find one star, BW I-155, that is a spectroscopic double, and another star, BW IV-203, with emission in the wings of its $\text{H}\alpha$ line.

Equivalent widths for nearly 2500 features have been measured, and abundance analysis of 11 BW bulge giants and two disk giants (α Boo and μ Leo) has been performed using LTE model atmospheres and a LTE spectrum synthesis code.

We have developed a simple yet useful equation for relating the infrared flux ratio R to the K -band bolometric correction. The infrared flux method has been used to determine accurate effective temperatures for 34 BW giants and shows that the $(J - K)$ temperature relation is the same for Baade's window and solar neighborhood red giants. The $(V - K)$ temperature relation in Baade's window, using photographic Arp (1965) V -band photometry, is cooler than the solar neighborhood relation by 115 K. This shift cannot be explained by an error in the adopted reddening, but could be due to a zero-point error in the Arp (1965) V -band photometry of 0.13 mag. This is entirely consistent with the observations of Blanco & Blanco (1984), who show that for BW stars in the range $V \sim 15.0$ – 16.0 , the Arp (1965) V photometry gives values that are, on average, greater than CCD V -band photometry by about 0.1 mag. Future study of Baade's window giants should include CCD V -band photometry to resolve the problem of the zero-point difference between $(V - K)$ and IR flux temperatures. Photometry from the U band to far-IR bands is recommended, to increase the pool of data for IP flux temperatures.

We identify a divergence between $(J - K)$ and $(V - K)$ color temperatures for the most metal-rich stars in our sample. We suppose that this difference may be due to blanketing in the V band, and might also be responsible for the difference between our final abundances and metallicities based on Washington photometry. For the abundance analyses of these stars we adopted the $(J - K)$ color temperatures; if we had included the $(V - K)$ color temperatures for these stars, the temperatures and $[\text{Fe}/\text{H}]$ values would have been reduced.

We present a list of continuum regions found in our spectrum of μ Leo; these will be of use to those interested in abundance analysis of super-metal-rich K giants.

To address the CN line blanketing affecting abundances de-

rived from the EW method, we have used spectrum synthesis to identify atomic lines which are clear of significant blending at the resolution of our spectra. Our spectrum syntheses cover 1460 Å in 0.01 Å steps and include more than 8000 ^{12}C N and ^{13}C N lines and 200 atomic lines at a variety of metallicities and CN strengths. We find that for μ Leo 75% of the originally selected Fe I lines are significantly affected by CN. Surprisingly, we find that, even at solar metallicity, lines originally thought to be clean are in fact significantly blended with CN. Using spectrum synthesis, we have categorized our original atomic line list into sublists of lines useful for abundance analysis at different metallicities. These lists of lines, relatively unaffected by CN blends, should be useful for future work on SMR and solar metallicity K giants.

We have compiled a set of atomic gf values, based on a critical investigation of the literature. In particular, for Fe I, we have found relations to transform gf values from a variety of sources onto the preferred system of O'Brian et al. (1991). In line with recent results, we adopt a low solar iron abundance, $\epsilon(\text{Fe}) = 7.52$, which is in excellent accord with the meteoritic value.

Our [Fe/H] results are lower than those found for the same stars by R88 (solution 1) by an average of 0.32 dex. We demonstrate that this difference may be partially due to the assumption by R88 that [Mg/Fe] = 0.00 in the bulge giants. However, we find that our [Fe/H] results are systematically lower than the R88 (solution 2) Fe-only index by 0.25 dex. The mean [Fe/H] for our 11 bulge giants is -0.33 dex.

When we use linear regression correlations to transform the R88 [Fe/H] scale, we find that the average corrected [Fe/H] for the R88 bulge sample is -0.25 dex for solution 1; for solution 2 we find an average of -0.21 dex. In addition, a comparison of the corrected R88 bulge with the solar neighborhood GK giant survey of McW90 shows that the mean of the two [Fe/H] distributions are identical, to within the measurement uncertainties. It does appear that the width of the corrected R88 [Fe/H] sample is broader than seen in the solar neighborhood. This can be understood by the lack of the so-called G dwarf problem in Baade's window. An important point is that the most metal-rich bulge star in our sample, BW IV-167 (which is the second highest in [Fe/H] of R88 solution 1), is not significantly more metal-rich than the SMR disk giant μ Leo.

The corrected R88 [Fe/H] distribution and our confirmation of the $(J - K)$ temperature scale limits the explanation of the FW87 IR H-R diagram to a luminosity discrepancy; the BW giants are more luminous than expected for globular cluster giants of the same temperature and metallicity. If the BW giant branch follows the 47 Tuc locus, as claimed, we find that this is consistent with a mean BW giant mass *no less* than about $1.10 M_{\odot}$; our formal mean mass of $1.35 M_{\odot}$ is unusually high and almost certainly in error. If the BW stars possess primordial [O/Fe] = 0.0, then the mean BW giant mass can be reduced by as much as $0.08 M_{\odot}$, to $1.27 M_{\odot}$.

Tentative confirmation of a low abundance scale for Baade's window is given by an ongoing spectroscopic analysis of a sample of M giants by Smith (1993). Our low [Fe/H] results are in conflict not only with the R88 spectroscopic results but also with the FW87 findings of superhigh metallicities suggested by the spectral type trend with $(J - K)$, and high CO indices seen in bulge M giants. The low-metallicity scale is also in conflict

with the high [Fe/H] values claimed by Geisler & Friel (1992) and Tyson (1991), which both relied upon Washington photometry of GK giants in BW.

Our mean [Fe/H] for the bulge suggests a zero Galactic [Fe/H] gradient within the solar circle. This conflicts with the conclusions of recent studies of the Galactic [Fe/H] gradient but does not conflict with the data presented in these studies. It is our opinion that the Galactic gradient studies have been affected by measurements outside the solar circle, where the metallicities do decrease markedly with Galactocentric distance; this may explain the considerable dispersion in published Galactic [Fe/H] gradients. Furthermore, theoretical modeling of Galactic metallicity evolution by Larson (1976) predicts a zero gradient within about 10 kpc of the Galactic center, and a relatively sharp decline outside this distance.

Detailed chemistry of the 11 stars shows that Mg and Ti are enhanced in almost all stars. The high Ti/Fe ratio can be used to explain qualitatively the observation of FW87 that the bulge M giants are too blue for their spectral type; we believe this is due to enhanced TiO band strengths rather than to an overall metallicity increase. This, combined with the sensitivity of the CO index in M giants to atmospheric parameters, is sufficient to show that the FW87 observations are consistent with a bulge of "normal" metallicity.

We point out that the high M giant frequency found in BW (Nassau & Blanco 1958; Blanco 1965) is easier to understand with an enhancement of Ti only rather than an overall metal enhancement. In the case of an enhancement of all metals, the TiO band strength increases owing to larger Ti (and to a lesser extent O) number density; but the increased metal abundances lead to higher electron number densities, and therefore a greater H^- continuum opacity, which works to diminish the increase in TiO band strength. In a situation where overall metal abundance remains largely unchanged, the continuum opacity is also unchanged, and the full effect resulting from an increase in TiO number density is reflected by the stronger TiO bands.

Our high Mg and Ti abundances indicate that most of the chemical enrichment in the bulge was performed by Type II SNe. However, our Si and Ca abundances seem to follow the relation seen in the Galactic disk and halo, contrary to the results for Mg and Ti. If our abundances are correct, we conclude that the currently popular simple picture of Galactic α -element enrichment is not tenable in the bulge, and stars responsible for the history of bulge nucleosynthesis differed from those of the disk and halo. Under the assumption of solar C/Fe and N/Fe ratios we have computed oxygen abundances for six of our stars, based on the [O I] line at 6300 Å. We find that the results may be described by [O/Fe] = 0.00, although BW I-012 is found to have [O/Fe] = 0.3. The unusually low oxygen abundances might be understood as a consequence of O burning, seen in many of the brightest globular cluster giants (Snedden et al. 1991; McWilliam et al. 1992).

At least one of our stars shows evidence of nonprimordial nuclear processing of the stellar envelope material, having an extreme enhancement of Li. It is remarkable that we should have encountered such a star in a sample of only 11 K giants, since the frequency among disk K giants is approximately 1% (Brown et al. 1989). Since the likely explanation for super-Li-rich K giants is binary mass transfer from an evolved compan-

ion, our detection of a super-Li-rich star could be the result of a high binary frequency in our sample of stars. The spectroscopic double BW I-155 supports this notion if it is in fact a binary system.

In addition to the super-Li-rich K giant, we find unusually high abundances for the elements Al, Sc, and Co in almost all of our stars. It is of great interest to determine whether the high abundances are due to binary evolution or some other stellar evolution mechanism, or were in fact unusually high in the primordial material. In the seemingly unlikely event that our sample of bulge K giants are predominantly mass transfer binaries, then these stars should all have white dwarf companions, rather like the barium stars. Possible white dwarf companions could be detected from radial velocity studies, low-resolution UV spectroscopy, and observations of the He I 10830 Å line (see Brown et al. 1990). Without doubt it is necessary to search for these high abundances in a larger sample of bulge K giants, especially those of lower luminosity. If solar metallicity bulge giants with high [Al/Mg] are the norm, then the trend of [Al/Mg] predictions by Arnett (1971) and Woosley & Weaver (1982) (confirmed observationally in the disk) must be revised for the bulge.

For V, Cr, Ni, and probably Na we find that the abundances relative to iron behave as seen in the disk and halo systems. We also find that our *s*-process index, utilizing Y and La lines, scales with the iron abundance in our stars, just as is observed in disk stars (Lambert 1987).

Taken at face value, the Eu (an *r*-process element) abundances for two of our most metal-rich bulge giants lie above the mean for solar neighborhood giants, and are in qualitative agreement with nucleosynthesis dominated by Type II supernovae. However, we give this result with some caution, and suggest that additional *r*-process lines (from Eu and other elements like Dy) be studied in these stars with greater spectral resolution and S/N to test this result.

Future abundance studies of bulge stars should attempt to confirm the abundance patterns found in this work using spectra of higher S/N and resolution, to facilitate measurement of weaker lines, which are less sensitive to upper atmosphere and microturbulent velocity effects.

The abundance trend of [S/Fe] with [Fe/H] would be interesting to measure because it should mimic the run of [Ca/Fe] and [Si/Fe] to [Fe/H], as Si, S, and Ca are thought to be produced in the same nuclear burning processes. Mg abundances could be checked using the MgH lines near 5100 Å.

Abundances for the CNO group of elements must be derived simultaneously; this would be facilitated by high-S/N, high-re-

solution spectra of C₂ features near 5100 Å, the [O I] line at 6300 Å, and CN red system lines. CNO and Al abundances of lower luminosity bulge giants would be of particular interest, since such stars are not expected to have altered their atmospheric O or Al abundances.

To confirm the high abundance of Ti in the bulge, accurate Ti oscillator strengths for weak lines will be necessary. We note that the Ti lines are so strong that lines of 0.5 mÅ in the solar spectrum are nearly 100 mÅ in the bulge K giants. This presents a significant problem for solar *g f* values of lines less than 100 mÅ in the bulge giants, and it may be necessary in some cases to resort to *g f* values determined from standard giants, such as α Boo. We remind the workers in atomic physics that there is a great need for laboratory oscillator strengths of weak lines in the red region of the spectrum, particularly Mg, Si, Ca, Ti, and Fe.

It is of interest to understand the discrepancy between our results and the Washington photometry of Geisler & Friel (1992) and Tyson (1991); high-quality spectra of their most metal-rich stars would be of great use in this regard.

Finally, in the course of this investigation it has become apparent to us that the spread in abundances derived in high-resolution analyses is in part due to blends, not only with molecular lines but also with lines due to other atomic species. Therefore, significant improvement in the accuracy of abundance results might be achieved with the detailed consideration of potential blends for each line, from all sources. In this matter the line list of Kurucz (1992) is especially helpful.

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